Doctoral Thesis

Coronal evolution of solar-like stars
X-ray spectroscopy of stars in star-forming regions and the solar neighborhood

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Abstract

Solar-like stars are strong X-ray emitters in both their pre-main sequence (PMS) and main-sequence (MS) phases. In analogy to the Sun, X-rays are thought to originate in a corona. However, in the case of pre-main sequence stars, accretion processes might influence the X-ray properties of the stars.

In this thesis, results from X-ray spectroscopy of main-sequence solar analogs, pre-main sequence solar-like stars and a Herbig Ae/Be star are presented and discussed. All X-ray spectra have been obtained by the Reflection Grating Spectrometers (RGS) and the European Photon Imaging Cameras (EPIC) on board the XMM-Newton satellite.

In the first part of the thesis, high-resolution (RGS) X-ray spectra of a sample of six main-sequence G-type stars with ages between \(0.1\) Gyr and \(1.6\) Gyr have been analyzed. Using individual spectral lines, the Emission Measure Distributions (EMD) and the coronal abundances have been derived. As a solar analog evolves, its rotation rate decreases and its internal magnetic dynamo weakens, resulting in a decrease of magnetic activity and a decrease of the star’s X-ray luminosity. The mean coronal temperatures derived from the EMDs decrease from \(10\) MK for the youngest stars to \(4\) MK for the oldest star in our stellar sample. These results have been interpreted with a model in which the coronal emission is produced by a superposition of stochastically occurring flares; more active stars are found to require a larger range of flare energies than less active stars. Abundances change from an inverse First Ionization Potential (FIP) effect, where abundances with high FIP are enhanced with respect to abundances with low FIP, to a solar-like FIP effect at ages \(0.3\) Gyr.

The analysis has then been extended to pre-main sequence stars in the Taurus-Auriga complex. The results presented here are part of a large survey, the “XMM-Newton Extended Survey of the Taurus Molecular Cloud” (XEST). High- and medium-resolution spectroscopy has been used to study the differences in the X-ray output of low-mass accreting stars (classical T Tauri Stars, CTTS) and non-accreting stars (weak-line T Tauri stars, WTTS). For nine PMS stars, high-resolution RGS spectra were obtained. For two accreting stars, BP Tau and AB Aur, the electron densities were derived from the \text{O}{\textsc{vii}} triplets. Using the \text{O}{\textsc{vii}}/\text{O}{\textsc{viii}} flux ratio to quantify the cool plasma, a soft excess has been discovered; this feature seems to be a general feature in the
spectra of CTTS, but is not present in WTTS and zero-age main-sequence stars. In BP Tau, the O\textsubscript{VII} triplet suggests high electron densities for this cool plasma ($n_e = 3.4 \times 10^{11}$ cm\textsuperscript{-3}), a signature that at least part of the X-ray emission might originate in accretion shocks. However, low electron densities ($n_e < 10^{10}$ cm\textsuperscript{-3}) are inferred for the accreting Herbig star AB Aur. No significant, systematic differences are present between the abundances of accreting and non-accreting stars. Rather, the Ne/Fe coronal abundance ratio is a function of spectral type. It systematically increases from G to early M stars.

In a second approach, results from the medium-resolution EPIC spectra have been used to study systematics in the X-ray properties of CTTS and WTTS. Both CTTS and WTTS are found to be saturated in X-rays, i.e., the X-ray luminosity ($L_X$) is proportional to the stellar bolometric luminosity ($L_\star$), but the saturation level is different: for CTTS, $L_X = 10^{-3.73 \pm 0.05} L_\star$ while for WTTS, $L_X = 10^{-3.39 \pm 0.06} L_\star$. Furthermore, WTTS are more X-ray luminous than CTTS by a factor of $\approx 2$. A correlation between the average coronal temperature, $T_{\text{av}}$, and $L_X$ is present in WTTS but not in CTTS; overall, CTTS coronae are hotter than those of WTTS for low $L_X$. The X-ray emission of WTTS is found to be consistent with the coronal emission of main-sequence stars, whereas CTTS show a soft excess in high-resolution spectra, a lower X-ray activity, and generally hotter coronae in medium-resolution spectra. A model is proposed in which accreting material penetrates magnetic structures and cools them to temperatures that are no longer accessible by X-ray CCD detectors. This would reduce the $L_X$ determined from CCD spectra while would produce a soft excess that dominates the long-wavelength line spectra in grating spectrometers.

The high-resolution spectrum of the Herbig Ae/Be star AB Aur has been studied separately. The X-ray generation in Herbig stars is so far unexplained. AB Aur displays a soft spectrum characteristic of plasma at 1–6 MK. The X-ray light curve displays variability with a period of $\approx 42$ hours. AB Aur’s X-rays could be generated in a corona or in a magnetically confined wind. The latter alternative is supported by periodicities of optical lines formed in the wind, whose periods are consistent with the X-ray period.

Initial results from five \textit{XMM-Newton} observations of the Chamaeleon I star-forming region are presented. The X-ray properties of CTTS and WTTS are analogous to those reported from the Taurus Molecular Cloud, except that both CTTS and WTTS indicate a similar X-ray saturation level ($L_X \approx 10^{-3.3} L_\star$). Comparing the hydrogen column densities ($N_H$) found from the spectral fits with the corresponding infrared extinctions ($A_J$), $N_H/A_J$ ratios significantly below the galactic values are found. This suggests a larger grain size of the dust in the vicinity of the Chamaeleon I stars.
Riassunto

Evoluzione della corona di stelle simili al Sole: spettroscopia a raggi X di stelle in regioni di formazione stellare e nei dintorni del Sole

Le stelle di pre-sequenza o di sequenza principale simili al Sole sono intense sorgenti di raggi X. Si ritiene che i raggi X siano prodotti nella corona di queste stelle, in analogia con quanto osservato nel Sole. Tuttavia, nelle stelle di pre-sequenza, i processi di accrescimento possono influenzare l’emissione di raggi X.

In questa tesi vengono presentati i risultati ricavati dallo studio di spettri a raggi X di stelle di sequenza principale analoghe al Sole, stelle di pre-sequenza simili al Sole e una stella Herbig Ae/Be. Gli spettri X sono stati ottenuti con gli strumenti “Reflection Grating Spectrometers” (RGS) e “European Photon Imaging Cameras” (EPIC) che si trovano a bordo del satellite *XMM-Newton*.

Nella prima parte di questa tesi, sono stati analizzati gli spettri ad alta risoluzione di sei stelle di sequenza principale appartenenti alla classe spettrale G. Tali stelle hanno età differenti comprese tra ≈ 0.1 miliardi di anni (Ga) e ≈ 1.6 Ga. Utilizzando il flusso delle diverse linee spettrali, sono state ricavate la Distribuzione di Emissione di Massa (DEM) e le abbondanze degli elementi presenti nelle corone. Mentre una stella invecchia, la velocità di rotazione diminuisce, così che la dinamo interna viene inibita, causando una diminuzione dell’attività magnetica e della luminosità X. La temperatura della corona, ottenuta dalle DEM, diminuisce da ≈ 10 MK per le stelle più giovani a ≈ 4 MK per le stelle più evolute presentate in questo studio. Questi risultati sono stati interpretati con un modello in cui l’emissione della corona è costituita da una sovrapposizione di brillamenti (“flares” in inglese) distribuiti stocasticamente. Secondo questo modello, le stelle più attive necessitano brillamenti con una gamma di energie più ampia. Le abbondanze degli elementi cambiano da una distribuzione in cui gli elementi con un Primo Potenziale di Ionizzazione (PPI) più elevato sono più abbondanti degli elementi con un minore PPI, ad una distribuzione simile a quella trovata nella corona del Sole, dove gli elementi con un basso PPI sono più abbondanti degli elementi con un PPI più alto. Questo cambiamento avviene ad un’età stellare di ≈ 0.3 Ga.

Abbiamo successivamente esteso la nostra ricerca a stelle di pre-sequenza che si trovano nella regione del Toro-Auriga. I risultati qui presentati fanno
parte di un progetto più esteso chiamato “XMM-Newton Extended Survey of the Taurus Molecular Cloud” (XEST). Utilizzando la spettroscopia a raggi X ad alta e media risoluzione, sono state studiate le differenze nella produzione di raggi X tra stelle che accrescono (solitamente chiamate CTTS) e stelle che non accrescono (WTTS). Per nove stelle di pre-sequenza sono stati ottenuti spettri ad alta risoluzione. Per due CTTS è stato possibile misurare la densità degli elettroni utilizzando le tre linee spettrali dell’ O\text{vii}. Utilizzando il rapporto tra i flussi misurati nelle linee dell’ O\text{vii} con quelli misurati nelle linee dell’ O\text{viii} è stato possibile determinare la quantità di plasma che si trova a temperature relativamente basse. Un eccesso di plasma a temperature basse è stato rilevato in tutte le CTTS, ma non nelle WTTS e nelle stelle appena entrate nella sequenza principale. Le tre linee dell’ O\text{vii} nello spettro di BP Tau suggeriscono una densità degli elettroni particolarmente alta ($n_e = 3.4 \times 10^{11} \text{ cm}^{-3}$) mentre le stesse linee nello spettro della stella Herbig AB Aur sono coerenti con una bassa densità ($n_e < 10^{10} \text{ cm}^{-3}$). Non sono state trovate differenze significative tra le abbondanze degli elementi delle CTTS e delle WTTS. Inoltre l’abbondanza del Ne è risultata 4-6 volte superiore all’abbondanza del Fe in tutte le stelle della classe spettrale K e M, mentre l’abbondanza del Fe è maggiore nelle stelle di tipo G.

Sono in seguito state analizzate le proprietà delle stelle che accrescono e delle stelle che non accrescono utilizzando spettri a media risoluzione ottenuti con EPIC. In entrambi i tipi di stelle la luminosità X ($L_X$) è risultata proporzionale alla luminosità bolometrica della stella ($L_\ast$), indicando una saturazione dell’emissione X. I livelli di saturazione sono però diversi: per le CTTS $L_X = 10^{-3.73\pm0.05} L_{\text{bol}}$, mentre per le WTTS $L_X = 10^{-3.39\pm0.06} L_{\text{bol}}$. Le WTTS risultano inoltre essere un fattore $\approx 2$ più luminose a raggi X della CTTS. È stata rilevata una correlazione tra la temperatura media delle corone ($T_{\text{eq}}$) e $L_X$ nelle WTTS ma non nelle CTTS; le CTTS risultano avere temperature più elevate anche quando hanno luminosità X basse. In conclusione, l’emissione osservata delle WTTS è coerente con l’emissione osservata nelle corone di stelle giovani di sequenza principale. D’altra parte, l’emissione osservata delle CTTS rivela un eccesso di plasma a basse temperature negli spettri ad alta risoluzione, mentre le temperature sono in media più alte e le luminosità in media più basse negli spettri a media risoluzione. È possibile spiegare questi risultati utilizzando un modello in cui il materiale di accrescimento penetra nelle strutture magnetiche della corona e la raffredda a temperature che non possono più essere misurate con i detettori CCD di EPIC. In questo caso, la luminosità $L_X$ diminuirebbe e verrebbe prodotto un eccesso di emissione a temperature più basse che potrebbe venir misurato solo in spettri ad alta risoluzione, in cui è possibile misurare le linee spettrali che si formano a tali temperature.

Lo spettro ad alta risoluzione della stella Herbig Ae/Be AB Aur è presentato separatamente. La produzione di raggi X in questo tipo di stelle rimane
tuttora inspiegata. Lo spettro di AB Aur rivela plasma relativamente freddo (1-6 MK), mentre la curva di luce mostra variabilità con un periodo di \( \approx 42 \) ore. L’emissione X può essere stata prodotta in una corona, oppure in venti stellari confinati dai campi magnetici. Quest’ultima alternativa è supportata dai periodi misurati in alcune linee dello spettro ottico e prodotte dai venti stellari: tali periodi sono infatti coerenti con il periodo misurato nella curve di luce X.

Nel capitolo 6 sono presentati i risultati preliminari di uno studio, composto da 5 osservazioni ottenute con XMM-Newton, della regione di formazione stellare del Camaleonte I. Le proprietà dell’emissione X di CTTS e WTTS sono risultate simili a quelle ottenute nel Toro, tranne che in questo caso le CTTS e WTTS si trovano allo stesso livello di saturazione dell’emissione X (\( L_X \approx 10^{-3.3} L_\odot \)). Confrontando l’assorbimento fotoelettrico (espresso in densità di colonna dell’idrogeno, \( N_H \)) misurato negli spettri con l’estinzione nell’infrarosso (\( A_J \)) corrispondente, sono stati trovati valori di \( N_H/A_J \) parecchio inferiori a quelli galattici. Ciò potrebbe suggerire che nel Camaleonte I si trovano granelli di polvere stellare di dimensioni superiori rispetto alla norma.
Riassunto
Chapter 1

Introduction

X-ray emission from our star, the Sun, was discovered about half a century ago. This high-energy radiation is emitted from plasma confined in magnetic structures in the Sun’s outer atmosphere that is continuously heated to temperatures above one million degrees. This outer, hot atmosphere is called the corona.

Radiation in the X-ray wavelength domain is not unique to the Sun. X-ray emission has been observed in a large fraction of Main-Sequence (MS) and Pre-Main Sequence (PMS) stars. However, as we will describe below, the X-ray luminosities and the X-ray temperatures of PMS stars and active MS stars may exceed solar values by several orders of magnitude.

The goal of this work is to study the coronal evolution of the Sun in time from the pre-main sequence phase, when it was still accreting material from the circumstellar disk, to an evolved phase on the main-sequence, corresponding to the present-day age. This work is motivated by the following question: How do the physical properties of the magnetic corona of a solar-like star on the main-sequence evolve in time? What are the characteristics of X-rays emitted in pre-main sequence solar-like stars? Are they also originating in a corona? Does accretion influence the X-ray output of pre-main sequence stars?

In this thesis we present results from the analysis of X-ray spectra of MS and PMS stars. In Chapter 2 we study high-resolution spectra of a sample of six solar analog stars, i.e., main-sequence stars of spectral type G with masses \( M \) of \( \approx 1M_\odot \). These six stars have different ages: The two youngest stars, 47 Cas B and EK Dra, are zero-age main-sequence (ZAMS) stars at the age of \( \approx 0.1 \) Gyr (i.e. they have just started hydrogen burning in their interiors and have therefore just settled on the MS). In contrast, the oldest star, \( \beta \) Com, is 1.6 Gr old. Chapters 3–5 continue these studies back to the pre-main sequence phase. Chapters 3–5 report results from the XMM-Newton Extended Survey of the Taurus Molecular Cloud (XEST): In Chapter 3, results from high-resolution spectroscopy of 9 PMS stars are reported, while in Chapter 4 the role of accretion in X-ray emission is discussed. The high-resolution X-
ray spectrum of the Herbig star AB Aur is presented in Chapter 5. Finally, in Chapter 6 we present initial results from an XMM-Newton survey of the Chamaeleon I star forming region. Chapters 2, 3, 4, 5, and 6 are based on Telleschi et al. (2005, 2007a,b,c, 2006).

This introduction is structured as follows: In Sect. 1.1, a short history of stellar X-ray astronomy is presented. The XMM-Newton observatory, used to obtain the observations on which this thesis is based, is described in Sect 1.2. Sect. 1.3 gives a brief overview of magnetic activity observed on the Sun. Stellar X-ray spectra and diagnostics of temperature, density and abundances are described in Sect. 1.4, while Sect. 1.5 is a concise overview of results obtained from previous X-ray observations of stars. Finally, the XEST project is introduced in Sect 1.6.

1.1 A Short History of Stellar X-Ray Astronomy

The field of X-ray astronomy was born only a few decades ago. X-rays themselves were discovered in 1895 by the German physicist Wilhelm Conrad Röntgen, an achievement for which he was awarded the first Nobel prize in 1901. Since X-rays are absorbed by the Earth’s atmosphere, astronomers can observe the sky at these wavelengths only by using rockets or satellites.

The study of stellar coronae started with observations of the solar corona. The presence of a hot gas surrounding the Sun was first indirectly inferred in 1939 from optical coronal lines of highly ionized species (Grotrian 1939; Edlen 1942). Direct detection of X-ray photons from the Sun succeeded 10 years later, during a rocket flight (Burnight 1949). In the early seventies, the Apollo Telescope Mount aboard the Skylab provided high-quality X-ray images of the full Sun, which together with data from previous rocket flights formed the basis of our modern knowledge of the solar corona (and thus stellar coronae).

The history of stellar X-ray astronomy started in 1975, when Catura et al. (1975) reported the detection of X-rays from Capella. The X-ray luminosity of this binary star was found to be $\approx 10^{31}$ erg s$^{-1}$, i.e. four orders of magnitude above the Sun’s X-ray luminosity, and the plasma electron temperature amounted to $8 \times 10^6$ K, also several times higher than temperatures commonly observed on the Sun. These results were confirmed by Mewe et al. (1975), who first proposed that the X-ray emission could arise from an enhanced solar-like corona. In the same year, the first stellar X-ray flares were observed on YZ CMi and UV Cet (Heise et al. 1975).

Haisch et al. (1977) discovered high-energy radiation from a MS star, Proxima Centauri, in the extreme ultraviolet range. In the following year, $\alpha$ Cen was identified as a low-activity coronal source, with an X-ray luminosity com-
parable to the Sun’s (Nugent & Garmire 1978).

The Einstein observatory, a satellite launched in 1978, brought a revolution in the field. A large number of stars were detected, and X-ray emission was recognized as a characteristic feature common to a wide range of stars covering essentially the whole Hertzsprung-Russell diagram. Furthermore, X-ray emission from young stars residing in star-forming regions was detected for the first time (e.g., Feigelson & DeCampli 1981). Feigelson & DeCampli (1981) reported the observation of 28 young, low-mass stars (T Tauri stars, named after the prototypical young star T Tau) in the Taurus Molecular Cloud and in the Orion Nebula. They found X-ray luminosities of \( L_X \approx 10^{30} \text{ to } 10^{31} \text{ erg s}^{-1} \) and reported the occurrence of large flares also in this class of stars. With the improved spectral quality of the Einstein observatory, multi-temperature and variable-abundance spectral fits became possible. This allowed to discover the co-existence of a cool and a hot plasma component in RS CVn binaries. Grating spectroscopy was introduced, allowing spectral lines or line blends to be resolved. However, the sensitivity was insufficient to derive precise information on the thermal structure of the coronae (conventionally expressed in terms of the emission measure distribution, EMD, see Sect. 1.4).

In June 1990 the ROSAT satellite was launched. This observatory was several times more sensitive than Einstein. The ROSAT All-Sky survey identified more than 100000 sources, including tens of thousands of coronal sources.

Today, stellar X-ray observations are predominantly conducted by two large observatories: Chandra and XMM-Newton. NASA’s Chandra X-ray Observatory was launched in July 1999, and ESA’s XMM-Newton followed only a few months later, in December 1999. From both observatories, high-resolution X-ray spectroscopy is available, permitting precise measurements of spectral lines to derive information on the emission measure distribution, the plasma composition, and electron densities of stellar coronae. The two observatories are complementary, XMM-Newton being more sensitive and Chandra offering higher spatial and spectral resolution. In this thesis, we present results based on data obtained from the XMM-Newton observatory.

### 1.2 The XMM-Newton Observatory

The European Space Agency’s X-ray Multi-Mirror Mission (XMM-Newton, Fig. 1.1) was launched on December 10, 1999. It is the largest astronomy satellite built and operated by ESA. It orbits the Earth in 48 hours in a highly eccentric orbit, allowing approximately 40 hours of contiguous observations to be recorded. A picture of the XMM-Newton payload is shown in Fig. 1.2 (from Bagnasco 1999). Four telescopes operate on the satellite: three co-aligned X-ray telescopes that have a focal length of 7.5 meters each, and a co-aligned optical telescope (optical monitor, OM). The optics of each X-ray
telescope consists of 58 gold-coated, nested Wolter-I-type mirrors. Each mirror is shaped to a paraboloid surface in front and hyperboloid surface at the rear, leading to double reflection of the incident X-rays. The mirrors provide a large effective area, making XMM-Newton the most sensitive X-ray telescope so far.

XMM-Newton carries three scientific instrument groups: the European Photon Imaging Cameras (EPIC), the Reflection Grating Spectrometers (RGS), and the Optical Monitor (OM). The EPIC instrument consists of three charge-coupled device (CCD) cameras, of which two are MOS (Metal Oxide Semiconductor) detectors and one is a PN detector. Both types of detectors provide imaging and medium-resolution spectroscopy. To obtain high spectral resolution, the RGS instruments are used. Grating structures are mounted on two telescopes, reflecting and dispersing about half of the incoming X-rays to a secondary focus, where they are collected by a CCD array. Five X-ray instruments are thus available at the same time. The OM allows simultaneous collection of optical/UV radiation.

1.2.1 The EPIC Cameras

The three EPIC cameras perform sensitive imaging observations over a circular field of view of 30 arcmin diameter each. The PN camera (Strüder et al. 2001) provides most counts, because about half of the photons that enter the telescopes feeding the MOS cameras (Turner et al. 2001) are reflected by the gratings and are dispersed to the RGS detectors. The two MOS cameras are
1.2. The *XMM-Newton* Observatory

composed of 7 CDD chips each. They operate in the energy range between 0.2 and 12 keV. The MOS Point Spread Function (PSF) has a full-width-at-half-maximum (FWHM) of about 4 – 5″ and an half-energy-width (HEW) of 13 – 15″. The PN camera is composed of 12 CCD chips. The characteristics of these CCDs are a high detection efficiency up to 15 keV, a low noise level, and fast readout. The PN has a PSF FWHM slightly larger than the MOS, namely ≈ 6″. The spectral resolving power of all EPIC cameras is $E/\Delta E \approx 20 – 50$.

The three cameras allow for different modes of data acquisition. In the ‘full frame mode’, all pixels of all CDDs are read, and the complete field of view is covered. In the partial ‘small window mode’ of the MOS cameras, only the central part of the central CCD is read. In this mode, the central CCD can be read faster so that pile-up from bright sources can be avoided. The PN camera can also be operated in a ‘large window mode’, in which only half of the area of all 12 CCDs are read out, and in a ‘small window mode’, where only a small part of the CCD covering the on-axis position is read out. Further, three blocking filters (thick, medium, and thin) are available for the three cameras to suppress contamination by ultraviolet (UV), infrared (IR), and optical photons.
Chapter 1. Introduction

1.2.2 The Reflection Grating Spectrometers

The RGS is described by den Herder et al. (2001). In two out of the three telescopes on board XMM-Newton, about 50% of the light is dispersed using reflection gratings. The light strikes the grating with an angle $\alpha$ and is reflected at the angle $\beta$. The dispersion equation is then

$$d \cdot \cos \alpha - d \cdot \cos \beta = m \cdot \lambda \quad m = 1, 2, 3, \text{etc.}$$

(1.1)

$\lambda$ being the wavelength and $m$ being the refraction order. The resulting spectra cover the wavelength range from 5 to 35 Å. In each of the two spectrometers, one CCD chip failed early in the mission, leaving a gap between 10.6-13.8 Å in the RGS1 and between 20.0-24.1 Å in the RGS2. The first order ($m=1$) spectral resolution is approximately 60-70 mÅ at each wavelength, corresponding to $\lambda/\Delta \lambda \approx 300$ at $\lambda \approx 20$ Å.

Figure 1.3 shows the RGS spectrum of the young (0.1 Gyr) solar analog star 47 Cas B as an example (the spectrum is shown in flux units, and the RGS1 and RGS2 spectra have been coadded). Most of the stronger emission lines are separated.

1.3 Magnetic Activity on the Sun

The solar corona offers the closest laboratory for the study of stellar coronae. The magnetic processes observed in the Sun are fundamentally important for stellar astronomers to understand analogous activity observed in solar-like stars, i.e., MS stars with spectral types between F and M. These stars are believed to have an internal structure roughly comparable with that of the Sun.
1.3. Magnetic Activity on the Sun

Figure 1.4: Yohkoh images of the solar corona taken close to maximum activity (1991, left) and close to minimum activity (1995, right).

Magnetic activity can be observed in different layers of the solar atmosphere: In the solar corona, the magnetic fields play a fundamental role by providing the energy for coronal heating, by confining the hot plasma in bright loops, and by releasing energy explosively in flares. Just below, in the chromosphere, cooler and denser plasma supported by magnetic fields can be observed. Finally, magnetic fields are expressed, for example, through sunspots in the photosphere, the lowest layer of the atmosphere.

Chinese astrologers reported the presence of sunspots already around 800 B.C. In the 17th century, Galileo studied the Sun in detail for the first time using a telescope, and discovered that the surface is dotted with dark spots. In the 19th century, long-term observations of the Sun led to the discovery of the 11-year activity cycle: the number and latitude of the sunspots follow a regular pattern, which is repeated every 11 years. Subsequent studies revealed that the Sun’s magnetic field changes polarity after each 11 year cycle, so that a full magnetic cycle lasts 22 years.

Sunspots are regions of the photosphere characterized by a lower temperature and high magnetic flux. Sunspots are modeled as the counterparts of the magnetic flux tubes in the convection zone that are wound up by rotation.

The outer layer of the solar atmosphere, the corona, is entirely controlled by magnetic activity. While the temperature of the photosphere is only 5800 K, the corona reaches temperatures exceeding $10^6$ K (1 MK). The extreme temperatures of the corona produce radiation predominantly in the UV, extreme UV (EUV), and X-ray regimes. The X-ray activity of the Sun is strongly modulated (following the 11 year cycle), as can be seen in the two soft X-ray images taken with the Yohkoh satellite in Figure 1.4. In the left image, taken
close to maximum activity, bright, loop-shaped magnetic structures are seen in which the plasma is confined. In the right image, taken close to minimum solar activity, almost no bright feature is visible. Peres et al. (2000) found that the total solar luminosity in the 0.1-3 keV band changes between $\approx 5 \times 10^{27} \text{erg s}^{-1}$ at solar maximum and $\approx 3 \times 10^{26} \text{erg s}^{-1}$ at solar minimum.

The Sun is the only star for which the structure and topology of the X-ray corona can be spatially resolved. Regions where sunspots are seen in visible light are bright in the X-rays. Here, the plasma is confined in close magnetic structures ("magnetic loops"). An example can be seen in Figure 1.5. The figure was taken by the TRACE satellite using a 171 Å bandpass filter that covers the emission line of Fe IX and is sensitive to plasma at $\approx 1 \text{MK}$.

The loops such as those seen in Figure 1.5 reflect the shape of the underlying magnetic fields. They evolve and change their appearance continually. Explosive phenomena, called flares, are observed. They occur in active regions when magnetic reconnection take place in a discontinuity between magnetic regions with antiparallel field components. The flare process thus reduces the complexity of magnetic field structures by liberating non-potential magnetic energy and transforming it into kinetic energy of particles and heat.

The "steady" heating mechanism of the corona is still debated, and different mechanisms have been proposed. One possibility is that the energy is released through magnetic reconnection events as observed in flares. However, large flares are observed only in magnetic active regions, while the corona remains hot also when active regions are not present (such as during the solar minimum). Parker (1988) proposed that the corona is heated by the collective contribution of very small flares called micro- or nano-flares. The energy input by each of these small flares is tiny, and it is therefore difficult to observe them individually.
1.4 X-ray Spectra

X-ray emission from a stellar corona is emitted by a hot, tenuous, and highly ionized gas. Temperatures usually exceed one million degrees. The tenuous gas is to a good approximation optically thin, such that the observed spectrum faithfully represents the microscopic emissions that occur in the plasma. X-ray spectra, like the one shown in Figure 1.3, are basically the sum of two emission processes: continuum emission and line emission.

The continuum emission in the temperature range relevant for our spectra is mainly produced by bremsstrahlung. A free electron interacts with an ion by Coulomb forces, described as a free-free (ff) transition between two continuum states of the ion. Two other processes contribute to the continuum: the free-bound (fb) transitions (by capture of the electron into a bound state of the ion and emission of a photon) and the two-photon processes \( (2\gamma) \), excitation of a metastable level in a H- or He-like ion followed by simultaneous emission of two photons.

Up to temperatures of \( \approx 10 \) MK, the X-ray spectrum is dominated by line emission. Line emission occurs when partly ionized atoms are excited by electron collisions. In particular, in the wavelength ranges observed with XMM-Newton (i.e., between 1 and 35 Å), a set of differently ionized Fe lines can be recorded (Fe xvi-Fe xxvi), together with lines of C, N, O, Ne, Mg, Si, and S. The most relevant lines for spectroscopy with XMM-Newton are described in Table 1.1. The table is, of course, not complete. For a complete list of lines we refer the reader to the ATOMDB home page (http://cxc.harvard.edu/atomdb/). Not all lines listed in the table are separated: for example, the Si xiii He\( \alpha \) triplet is not resolved by the RGS while the Ne ix He\( \alpha \) is blended with Fe xix lines. In this thesis, we measured the Ly\( \alpha \) lines of Mg, Si, and S and the He\( \alpha \) lines of Mg, Si, S, and Fe in EPIC spectra rather than in RGS spectra: Fe He\( \alpha \) and S transitions are outside the wavelength range accessible by the RGS, and Mg and Si lines are often very weak in the RGS, due to the low sensitivity of the instrument at these wavelengths. Because the He\( \alpha \) triplets of Fe, Mg, Si, and S are unresolved in the EPIC spectrum, only the resonance (\( r \)) lines are listed in Table 2.5.

Fig. 1.6 shows the RGS1 XMM-Newton spectrum of Capella (taken from Audard et al. 2001) resulting from 52.3 ks of observing time taken in 2000. The data are shown with black crosses, while the optimal spectral fit based on thermal plasma models is plotted in red. The relevant lines of the spectrum
### Table 1.1: Spectral lines relevant for high-resolution spectroscopy with XMM-Newton

<table>
<thead>
<tr>
<th>Ion</th>
<th>Wavelength (Å)</th>
<th>Transition</th>
<th>$\log T_{\text{max}}^a$ (K)</th>
</tr>
</thead>
<tbody>
<tr>
<td>Fe xxv$^b$</td>
<td>1.85</td>
<td>$1s^22p^1P_1 \rightarrow 1s^21S_0$</td>
<td>He$\alpha$ r 7.8</td>
</tr>
<tr>
<td>Fe xxiv</td>
<td>10.62</td>
<td>$1s^23p^2P_{3/2} \rightarrow 1s^22s^2^2S_{1/2}$</td>
<td>7.3</td>
</tr>
<tr>
<td>Fe xxi</td>
<td>11.74</td>
<td>$1s^22s^3d^1D_2 \rightarrow 1s^22s2p^1P_1$</td>
<td>7.2</td>
</tr>
<tr>
<td>Fe xxi</td>
<td>12.29</td>
<td>$1s^22s^22p^3d^3D_1 \rightarrow 1s^22s^22p^3^3P_0$</td>
<td>7.0</td>
</tr>
<tr>
<td>Fe xx</td>
<td>12.83</td>
<td>$1s^22s^22p_{1/2}^1P_{3/2} \rightarrow 1s^22s^22p^3^3S_{3/2}$</td>
<td>7.0</td>
</tr>
<tr>
<td>Fe xix</td>
<td>13.52</td>
<td>$2s^22p^3(2D)3d^3D_3 \rightarrow 2s^22p^5^3P_2$</td>
<td>6.9</td>
</tr>
<tr>
<td>Fe xvii</td>
<td>14.20</td>
<td>$2s^22p^4(1D)3d^2D_{5/2} \rightarrow 2s^22p^5^2P_{3/2}$</td>
<td>6.8</td>
</tr>
<tr>
<td>Fe xvii</td>
<td>13.77</td>
<td>$2s^22p^6^1P_1 \rightarrow 2s^22p^6^1S_0$</td>
<td>6.8</td>
</tr>
<tr>
<td>Fe xvii</td>
<td>15.01</td>
<td>$2s^22p^5(2P)3d^1P_1 \rightarrow 2s^22p^6^1S_0$</td>
<td>6.7</td>
</tr>
<tr>
<td>Fe xvii</td>
<td>15.26</td>
<td>$2s^22p^5(2P)3d^3D_1 \rightarrow 2s^22p^6^1S_0$</td>
<td>6.7</td>
</tr>
<tr>
<td>Fe xvii</td>
<td>16.78</td>
<td>$2s^22p^5(2P)3s^1P_1 \rightarrow 2s^22p^6^1S_0$</td>
<td>6.7</td>
</tr>
<tr>
<td>Fe xvii</td>
<td>17.05</td>
<td>$2s^22p^5(2P)3s^3P_1 \rightarrow 2s^22p^6^1S_0$</td>
<td>6.7</td>
</tr>
<tr>
<td>Fe xvii</td>
<td>17.10</td>
<td>$2s^22p^5(2P)3s^3P_1 \rightarrow 2s^22p^6^1S_0$</td>
<td>6.7</td>
</tr>
<tr>
<td>O viii</td>
<td>18.97</td>
<td>$2p^2P_{3/2,1/2} \rightarrow 1s^22^2S_{1/2}$</td>
<td>Ly$\alpha$ 6.5</td>
</tr>
<tr>
<td>O viii</td>
<td>16.01</td>
<td>$3p^2P_{3/2,1/2} \rightarrow 1s^22^2S_{1/2}$</td>
<td>Ly$\beta$ 6.5</td>
</tr>
<tr>
<td>O viii</td>
<td>21.6</td>
<td>$1s2p^1P_1 \rightarrow 1s^21^1S_0$</td>
<td>He$\alpha$ r 6.3</td>
</tr>
<tr>
<td>O viii</td>
<td>21.8</td>
<td>$1s2p^3P_1 \rightarrow 1s^21^1S_0$</td>
<td>He$\alpha$ i 6.3</td>
</tr>
<tr>
<td>O viii</td>
<td>22.1</td>
<td>$1s2s^3S_1 \rightarrow 1s^21^1S_0$</td>
<td>He$\alpha$ f 6.3</td>
</tr>
<tr>
<td>Ne x</td>
<td>12.13</td>
<td>$2p^2P_{3/2,1/2} \rightarrow 1s^22^2S_{1/2}$</td>
<td>Ly$\alpha$ 6.8</td>
</tr>
<tr>
<td>Ne ix</td>
<td>13.45</td>
<td>$1s2p^1P_1 \rightarrow 1s^21^1S_0$</td>
<td>He$\alpha$ r 6.6</td>
</tr>
<tr>
<td>Ne ix</td>
<td>13.55</td>
<td>$1s2p^3P_1 \rightarrow 1s^21^1S_0$</td>
<td>He$\alpha$ i 6.6</td>
</tr>
<tr>
<td>Ne ix</td>
<td>13.70</td>
<td>$1s2s^3S_1 \rightarrow 1s^21^1S_0$</td>
<td>He$\alpha$ f 6.6</td>
</tr>
<tr>
<td>C vi</td>
<td>33.73</td>
<td>$2p^6P_{3/2,1/2} \rightarrow 1s^22^2S_{1/2}$</td>
<td>Ly$\alpha$ 6.1</td>
</tr>
<tr>
<td>C vi</td>
<td>28.48</td>
<td>$3p^2P_{3/2,1/2} \rightarrow 1s^22^2S_{1/2}$</td>
<td>Ly$\beta$ 6.2</td>
</tr>
<tr>
<td>N vii</td>
<td>24.78</td>
<td>$2p^2P_{3/2,1/2} \rightarrow 1s^22^2S_{1/2}$</td>
<td>Ly$\alpha$ 6.3</td>
</tr>
<tr>
<td>N vii</td>
<td>20.91</td>
<td>$3p^2P_{3/2,1/2} \rightarrow 1s^22^2S_{1/2}$</td>
<td>Ly$\beta$ 6.3</td>
</tr>
<tr>
<td>N vi</td>
<td>28.79</td>
<td>$3s2p^1P_1 \rightarrow 1s^22^2S_{1/2}$</td>
<td>He$\alpha$ r 6.3</td>
</tr>
<tr>
<td>Mg xii$^b$</td>
<td>8.42</td>
<td>$2p^2P_{3/2,1/2} \rightarrow 1s^22^2S_{1/2}$</td>
<td>Ly$\alpha$ 7.0</td>
</tr>
<tr>
<td>Mg xii$^b$</td>
<td>9.17</td>
<td>$1s2p^1P_1 \rightarrow 1s^21^1S_0$</td>
<td>He$\alpha$ r 6.8</td>
</tr>
<tr>
<td>Si xiv$^b$</td>
<td>6.18</td>
<td>$2p^2P_{3/2,1/2} \rightarrow 1s^22^2S_{1/2}$</td>
<td>Ly$\alpha$ 7.2</td>
</tr>
<tr>
<td>Si xiii$^b$</td>
<td>6.65</td>
<td>$1s2p^1P_1 \rightarrow 1s^21^1S_0$</td>
<td>He$\alpha$ r 7.0</td>
</tr>
<tr>
<td>S xvi$^b$</td>
<td>4.72</td>
<td>$2p^6P_{3/2,1/2} \rightarrow 1s^22^2S_{1/2}$</td>
<td>Ly$\alpha$ 7.2</td>
</tr>
<tr>
<td>S xvi$^b$</td>
<td>5.04</td>
<td>$1s2p^1P_1 \rightarrow 1s^21^1S_0$</td>
<td>He$\alpha$ r 7.2</td>
</tr>
</tbody>
</table>

$^a$ Maximum line formation temperature.

$^b$ Lines measurable in the EPIC spectrum.
Figure 1.6: XMM-Newton spectrum of Capella, taken from Audard et al. (2001).
are marked in the figure. The lines listed in Table 1.1 (except the lines that can be seen only with the EPIC instrument) are present, along with a number of weaker lines.

1.4.1 Temperature Diagnostics

The observed X-ray spectrum depends sensitively on the electron temperature. Transitions that we observe in line emission can occur only in certain temperature ranges because the formation of ions is strongly temperature dependent. The maximum line formation temperatures of different lines are given in Table 1.1. The characteristic electron temperature can then, for example, be diagnosed using ratios of two lines of different ions of the same element.

Some constraints on the electron temperature can be derived qualitatively from the spectrum. Typical spectral features that are observed in very hot plasma are: i) a well developed bremsstrahlung continuum; ii) lines from highly ionized heavy atoms, like Mg X and XII, and Fe XX-X XIV; and iii) a comparatively high flux ratio between H-like and He-like resonance lines such as O VIII/O VII, Ne X/Ne IX, but also ratios such as those of Fe XVIII/Fe XVII. Cool plasma, on the other hand, displays a very weak continuum and produces predominantly emission lines of Fe XVII, O VIII and O VII, N VII and N VI, and C VI and C V.

However, coronal plasma cannot be isothermal. A coronal spectrum is the superposition of spectra emitted by plasma at different temperatures. Conventionally, the thermal distribution is given by the differential emission measure distribution (DEM):

\[ Q(T) = \frac{n_e n_H dV}{d \ln T} \]  \hspace{1cm} (1.2)

where \( n_e \) and \( n_H \) are the electron and hydrogen densities, and \( dV \) is a differential volume element at temperature \( T \). \( Q(T) \) contains information on the emitting plasma temperature and on the density-weighted amount of plasma that emits the observed X-rays. \( Q(T) \) is directly related to the flux measured in a spectral line \( j \):

\[ F_j = \frac{1}{4 \pi d^2} \int A \varphi_j(T) Q(T) d \ln T, \]  \hspace{1cm} (1.3)

where \( \varphi_j(T) \) is the line power per unit of emission measure (also referred as “emissivity”) and \( A \) is the abundance of the element that emits the observed line, with respect to some standard abundances (usually taken to be solar photospheric abundances). Each \( \varphi_j(T) \) includes all information on these standard abundances and on the atomic physics of the transition.

In order to derive the thermal structure (i.e., the DEM) from the observed spectrum, Eq. 1.3 must be inverted. For real observations \( F_j \), this inversion is
an ill-conditioned problem with a number of different solutions. Moreover, the coronal abundances have to be derived simultaneously from the same spectrum. In order to constrain and solve the problem, some additional constraints are therefore imposed, like smoothness, positivity, or functional form of the DEM. Some of these constraints may not be physically founded. The different inversion techniques that are commonly used to derive the thermal structures are explained in Güdel (2004a).

If high-resolution spectra are available, the inversion method can be applied to single lines, and a discretized emission measure distribution (EMD, to distinguish it from the continuous DEM) can be derived. The advantage of this method is that the EMD and the abundances can be derived sequentially. In order to obtain an abundance-independent EMD, lines from ions of only one element are used (usually Fe, which displays a sufficient number of lines that cover the temperature range between log $T[K] =$ 6.5 and log $T[K] =$ 8). Alternatively, temperature-sensitive flux ratios of He-like to H-like lines can be used. The EMD is iteratively estimated until the flux measured in each line agrees with the predicted flux, given by

$$ F_C = \frac{1}{4\pi d^2} \sum_i EM(i)\varphi(i), \quad (1.4) $$

where the $EM(i)$ and $\varphi(i)$ denote the emission measure and the line emissivity in the temperature bin $i$. More details on the methodology are given in Chapter 2.

### 1.4.2 Element Abundance Diagnostics

Qualitatively, the element abundance ratios can be derived from the spectrum by comparing lines of different atoms that are formed at similar temperatures. However, for lines formed at different temperatures a cross-correlation between the thermal structure and the abundance will be inevitable. This problem can be solved if an abundance-independent EMD is first obtained (see Sect. 1.4.1) using Fe lines. The Fe abundance can then be found by requiring that the continuum (mainly formed by H and He) is well fitted. To obtain the abundances of the other elements, we compared the flux measured in their emission lines with the flux predicted from the EMD. The ratio of these two fluxes gives the abundance relative to Fe, with respect to the corresponding solar ratios.

### 1.4.3 Density Diagnostics

Electron densities can be measured using density-sensitive lines, like those in the He-like triplets. In the wavelength range covered by grating spectroscopy with XMM-Newton, the best-resolved and usually brightest triplet is the O VII
Chapter 1. Introduction

The NeIX triplet would be a valuable alternative, but it is strongly blended with FeXIX lines. The O VII triplet is sensitive to electron densities in the range between \( \approx 10^{10} \text{ cm}^{-3} \) and \( \approx 10^{12} \text{ cm}^{-3} \) (Gabriel & Jordan 1969). The example of Capella is shown in Fig 1.7 (from Ness et al. 2003a), while the energy-level diagram for the transitions is shown in Fig 1.8. The excited states \( 1s2p\,^1P_1, 1s2p\,^3P_{1,2}, \) and \( 1s2s\,^3S_1 \) decay to the ground state \( 1s^2\,^1S_0 \) through the resonance line (r), the intercombination line (i), and the forbidden line (f), respectively.

In the case of low densities, the \( 1s2s\,^3S_1 \) excited state decays radiatively to the ground state. However, in the case of high densities, the more frequent collisions trigger the excitation from the upper level of the forbidden transition to the upper level of the intercombination transitions. In this case, the flux measured in the f line decreases while the flux in the i line increases.

The measured ratio of the forbidden and intercombination lines, \( \mathcal{R} = f/i \), can be written as

\[
\mathcal{R} = \frac{\mathcal{R}_0}{1 + \phi/\phi_c + n_e/N_c}
\]

(Blumenthal et al. 1972), where \( \mathcal{R}_0 \) is the low-density limit (\( \approx 3.85 \) for O VII), \( n_e \) is the electron density, \( N_c \) is the critical density at which \( \mathcal{R} \) drops to \( \mathcal{R}_0/2 \), and \( \phi/\phi_c \) describes the influence of the external radiation field, \( \phi_c \), being the critical photoexcitation rate.

The influence of the external radiation field, \( \phi/\phi_c \), becomes relevant for
1.4. X-ray Spectra

stars with $T_{\text{eff}} \gtrsim 10^4$ K. According to Blumenthal et al. (1972), the influence of the radiation field is then given by

$$
\phi/\phi_c = \frac{3(1 + F)c^3}{8\pi\hbar\nu^3} \frac{A(1s2p^3P \rightarrow 1s2s^3S_1)}{A(1s2s^3S_1 \rightarrow 1s^21S_0)} \nu \nu, \quad (1.6)
$$

where the radiation field energy $\nu \nu$ is given by Planck’s equation

$$
\nu \nu = W \frac{8\pi\hbar\nu^3}{c^3} \frac{1}{exp(\hbar\nu/kT_{\text{eff}}) - 1}, \quad (1.7)
$$

in which the dilution factor $W$ was introduced (Mewe & Schrijver 1978):

$$
W = \frac{1}{2} \left( 1 - \left[ 1 - \left( \frac{R}{d} \right)^2 \right]^{1/2} \right), \quad (1.8)
$$

where $R$ is the stellar radius and $d$ is the distance of the source from the center of the star (see Ness et al. 2001 for further details). At the surface, $W = 1/2$. In Eq. 1.6, $A$ denotes the spontaneous transition probabilities ($8.12 \times 10^7$ s$^{-1}$ and $1.04 \times 10^3$ s$^{-1}$, respectively, for the O VII triplet according to Blumenthal et al. 1972), while $F$ is an expression that accounts for the collisional excitation rates and was evaluated at the maximum line formation temperature for O VII ($\approx 2$ MK), yielding $F = 0.42$ (Blumenthal et al. 1972).
1.5 X-Ray Observations of Stars

X-ray emission has been observed in almost all types of stars across the Hertzsprung-Russell diagram. The different types of stars presented in this thesis and their coronal properties are briefly described in this section.

1.5.1 Main-Sequence Stars

Stars of spectral types F to mid-M (grossly defining “solar-like” stars) are thought to have an internal structure similar to the Sun, which allows magnetic activity to be generated through the $\alpha\Omega$-dynamo mechanism. Theory predicts the absence of a magnetic dynamo in early A stars, B and O-type stars. These latter stars do not have an outer convection zone. The X-rays detected in stars with spectral type O or B are thought to be generated in shocks that develop in unstable winds. In contrast, at the cool end of the main sequence, stars with spectral type later than M5 become fully convective. The $\alpha\Omega$-dynamo can thus not work. However, recent studies of low-mass stars show that there are no significant changes between the X-ray properties of stars before $\approx$ M5 and stars with lower masses (see for example Delfosse et al. 1998). It is still unknown how the dynamo mechanism operates in these low-mass stars, and how their magnetic activity is generated. A possibility is that a turbulent dynamo mechanism (Giampapa et al. 1996) is at work, or that the central convection is, for unknown reasons, suppressed (allowing an $\alpha\Omega$-dynamo mechanism for the generation of the X-rays).

1.5.2 T Tauri Stars

Low-mass PMS stars are usually called T Tauri stars (TTS). These stars are divided into two groups according to the strength of their H\(\alpha\) lines: the classical T Tauri stars (CTTS) show strong H\(\alpha\) lines, typically with an equivalent width (EW[H\(\alpha\)]) > 10 \text{ Å}, a sign that the stars are accreting material from the circumstellar disk. In weak-lined T Tauri stars (WTTS), on the other hand, the H\(\alpha\) emission line fluxes are small, which is a sign that accretion has ceased. Alternatively, based on infrared (IR) observations, Young Stellar Objects (YSO) have been ordered in classes according to their infrared excess. Deeply embedded protostars at the start of their accretion phase are “Class 0” objects, more evolved protostars still embedded in their envelope are “Class I” objects, stars with a circumstellar disk are “Class II” objects, and stars with no IR excess are “Class III” objects. CTTS are roughly consistent with Class II objects, while WTTS are consistent with Class III objects. The difference between the two classifications is that the optical classification distinguishes accretors from non-accretors (based on the strength of the H\(\alpha\) lines), while the IR classification distinguishes stars with a circumstellar disk from stars with...
1.5. X-Ray Observations of Stars

no (thick) disk (as the IR excess measured in Class II stars arises from the thermal emission from the dust in the disk).

Both WTTS and CTTS exhibit X-ray variability on time scales of hours to days. Their X-ray emission is mostly consistent with a scaled-up version of the X-ray emission observed in main-sequence stars or the Sun. It has therefore been interpreted as arising from a corona.

However, the origin of X-ray emission in TTS is debated. While the bulk of the X-ray emission does seem to originate from a corona, young stellar objects, especially in their early evolutionary stage, are believed to be fully convective in their interior, and therefore the generation of magnetic fields through the αΩ-dynamo mechanism, which operates in the Sun, is not possible. Other types of dynamo have been proposed (see e.g., Giampapa et al. 1996). Moreover, in accreting stars, X-rays could also be generated in accretion shocks (Kastner et al. 2002) or, in jet-driving accreting stars, in shocks at the base of the jets (Bally et al. 2003; Güdel et al. 2005a, 2007c). We will further discuss accretion induced X-ray emission in Sect. 1.5.8.

1.5.3 Herbig Ae/Be Stars

Herbig Ae/Be stars are young, intermediate-mass (∼ 2 – 10M_☉) stars. They are accreting, pre-main sequence stars and are therefore considered to be the intermediate-mass analogs to CTTS. The nature of X-ray emission in such stars is controversial. Herbig stars, like MS B- and A-type stars, are believed to be fully radiative in their interiors. Models developed to explain their X-ray radiation include non-magnetic wind shocks, magnetically confined wind shocks, accretion shocks, coronal emission, or emission due to the interaction with the circumstellar disk.

X-ray emission from Herbig stars often displays flares and X-ray temperatures exceeding 10 MK (e.g., Hamaguchi et al. 2000; Skinner et al. 2004), favoring magnetic processes. The X-ray properties are then very similar to those observed in CTTS and, given the high rate of binaries and multiples found among Herbig stars (Feigelson et al. 2003), X-rays may in fact originate from an unresolved low-mass companion.

1.5.4 Rotation-Activity Relation

The X-ray activity in solar-like main sequence stars is strongly correlated with the stellar rotation period: stars that rotate faster exhibit higher X-ray luminosities (Pallavicini et al. 1981). Young MS stars rotate fast but slow down as they age due to the transport of angular momentum away from the star by a magnetic stellar wind. Therefore, X-ray activity is also anti-correlated with age. The dynamo efficiency depends also on the depth of the convective zone. The X-ray activity has therefore often been correlated with the Rossby number.
instead of the rotation period, where the Rossby number is defined as the ratio between the rotation period and the convective turnover time \( R_0 = P/\tau_c \).

X-ray activity is usually expressed as the ratio between the X-ray luminosity, \( L_X \), and the stellar bolometric luminosity, \( L_{\text{bol}} \). \( L_X/L_{\text{bol}} \) is a measure of the energy fraction required to heat the corona. The observed \( L_X/L_{\text{bol}} \) values span several orders of magnitude, from \( 10^{-7} \) to \( 10^{-3} \). The activity-rotation relation is then found to follow the relation

\[
\frac{L_X}{L_{\text{bol}}} \approx R_0^{-2}.
\] (1.9)

For fast rotators, this relation saturates at \( L_X/L_{\text{bol}} \approx 10^{-3} \) so that a dependency on the rotation period or the Rossby number is no longer present. Saturation occurs in all stellar classes, but the onset of saturation changes, depending on the spectral type. For solar analogs, saturation occurs at \( L_X/L_{\text{bol}} \gtrsim 1000 \) times the \( L_X/L_{\text{bol}} \) value for the non-flaring Sun. The age-activity relation, however, suggests that the young Sun was more active. The physical causes of saturation are still not understood.

1.5.5 Coronal Temperature-Activity Relation

A relation between the coronal temperature and activity has been found in X-ray observations of MS stars: Stars that display higher activity are found to have hotter coronae (see for example Schrijver et al. 1984; Güdel et al. 1997a). The causes of this relation are not clear. In active stars the corona is thought to be dominated by active regions, while the surface fraction occupied by quiet regions and coronal holes is small. With increasing activity, more extended, denser, and hotter active regions are responsible for both the higher X-ray luminosity (or activity) and the higher coronal temperature. This is analogous to conditions on the Sun where during solar minimum, a small fraction of the surface is covered by active regions, while this fraction is much larger during solar maximum (Orlando et al. 2001). However, even if the Sun were completely covered by active regions, the very high luminosities measured in ZAMS solar-like stars could not be reached (Orlando et al. 2001). As an alternative, extreme model, the observed coronal emission could be entirely generated by dynamic, flaring loops. Then, the relation between activity and temperature would be explained by more numerous interactions between adjacent magnetic loops in active stars (see Chapter 2).

1.5.6 Abundances

Many studies have indicated that the abundances in a stellar corona are different from those in the photosphere. This effect has first been observed in the Sun, where the ratio between coronal and photospheric abundances depends
1.5. X-Ray Observations of Stars

on the First Ionization Potential (FIP) of the elements. Elements with a FIP larger than $\approx 10$ eV (C, N, O, Ne, Ar) show abundance ratios with respect to hydrogen similar to those found in the solar photosphere, while elements with FIP $< 9$ eV (Si, Mg, Fe) are found to be overabundant (Meyer 1985a,b; Feldman 1992). This abundance anomaly has been referred to as the “FIP effect”. The abundance anomaly in the solar corona varies from feature to feature (Feldman 1992).

Studies of the coronal abundances in stars have suggested that anomalies are not only a solar characteristic. In inactive stellar coronae, a FIP effect similar to the Sun’s has been identified (Laming et al. 1996; Drake et al. 1997; Güdel et al. 2002). However, in very active stars, an opposite abundance pattern has been identified, with low-FIP elements being depleted with respect to the solar photosphere and underabundant with respect to high FIP elements (Brinkman et al. 2001; Güdel et al. 2001a). This effect was coined the “inverse FIP effect”. Such an effect was also observed in pre-main sequence stars (Imanishi et al. 2002). However, we emphasize that in most cases the stellar photospheric composition is unknown, and the coronal abundances are calculated relative to the solar photospheric abundances. It is commonly assumed that the stellar photospheric abundances are similar to those of the Sun. The present thesis will address this issue in more detail, comparing coronal abundances with stellar photospheric abundances (see Chapter 2 and Chapter 5).

Abundance anomalies are also observed during stellar flares: although the Fe abundance is usually depleted in active stars, it tends to increase during flares.

There is as of yet no consistent theory of these abundance anomalies. Current thinking is that a fractionation process occurs in the chromosphere, where low FIP elements are ionized while high FIP elements remain neutral. The process of fractionation probably involves electric fields, magnetic fields, gravitation and/or thermal diffusion.

Hénoux (1995) has presented a review of several sophisticated models that could explain the FIP effect observed on the Sun: Von Steiger & Geiss (1989) proposed a model where photoionization due to the UV radiation field occurs in the fractionation regions, and ionized species are subsequently trapped and transported to the corona. The model of Hénoux & Somov (1992) proposes the generation of electric currents along the magnetic flux tube. These currents and the orientation of the magnetic fields in the magnetic flux tube generate an ascending force that acts on ions but not on neutral particles.

A list of possible mechanisms to explain abundance anomalies in stars is also presented in Güdel (2004a). The listed models includes: a stratification of the atmosphere (Mewe et al. 1997), where different ions reside at different heights in the magnetic loops, depending on mass and charge; a model in which the coronal emission is due to a sequence of small flares that drive abundance
anomalies as sometimes observed in the Sun; for example, some solar flares are Ne rich. This may be because the high photoionization cross-section of Ne could make it behave like a low-FIP element under the influence of X-ray irradiation (Shemi 1991). Further, the increase of metallicity observed in large flares of active stars suggests that the coronal abundances increase during these events due to chromospheric/photospheric evaporation (Ottmann & Schmitt 1996; Mewe et al. 1997; Güdel et al. 1999). Another interesting model has been presented by Vilhu et al. (1993), who suggested that the depleted Fe abundance measured in active stars might be due to an enhanced continuum level manifested as an additional power-law bremsstrahlung component produced by the impact of non-thermal electrons in the chromosphere. Low metallicity is indeed found in stars that are strong non-thermal radio emitters (Güdel et al. 2002). During coronal evolution, around the time when the inverse-FIP effect changes to a FIP-effect, the non-thermal radio emission drops (Güdel et al. 2002). Non-thermal electrons in active stars may build up a downward-pointing electric field that could cause the depletion of low-FIP elements in the corona.

Another model that aimed at explaining both the FIP-effect and the inverse FIP-effect was proposed by Laming (2004). According to this model, the coronal loops are heated by waves, which penetrate the loop footpoint and heat the ions, but not the neutral atoms. Laming (2004) found that some waves might exert an upward force, resulting in a FIP effect. Varying the wave energy density, they found some cases where a downward force is present that could result in the observed inverse FIP-effect.

1.5.7 Flares

Flares are events of sudden energy release. According to present-day models, the energy is stored in non-potential magnetic fields, and a flare occurs when reconnection of anti-parallel magnetic fields takes place. This explosive energy release becomes measurable across the electromagnetic spectrum. In the corona, flares can heat the plasma up to tens of millions of degrees, emitting radiation that can be observed in the soft and hard X-ray domains. In Hα, flares appear as rapid brightenings in the chromosphere. Radio bursts are also observed during powerful flares, which indicates the presence of high-energy electrons in the coronal magnetic fields.

On the Sun, two major types of soft X-ray flares are observed: compact and long-duration ("two-ribbon") flares (Pallavicini et al. 1977). Compact flares are smaller and shorter events (lasting a few minutes) that are produced by a small number of loops, possibly by the interaction between neighboring loops. The loops remain nearly unchanged in shape and position during the event. In contrast, two-ribbon flares show much larger structures, and their decay time is longer, lasting 1-2 hours. Usually, they involve an "arcade" of magnetic
loops. They are most likely produced by an opening up of magnetic fields after a filament ejection and a subsequent relaxation by closing the field lines. During the flare, two ribbons of Hα emission form that move apart during the main phase of the flares (hence the “two-ribbon” designation).

In Figure 1.9, a schematic illustration of a two-ribbon flare is shown, as taken from Martens & Kuin (1989). Closed lines represent a helical flux tube that connects back to the photosphere. Reconnection takes place in the current sheet, where magnetic energy is converted into kinetic energy and heat. Hard X-rays are generated during flares when accelerated particles impact and heat chromospheric material at the loop footpoints. The soft X-rays are instead generated by the cooling of plasma that has evaporated in the post-flare loops. An example of a flare observed on the Sun with TRACE is shown in Fig.1.10.

Flares consistent with these two flare variants are observed in stars as well. However, extreme cases have also been recorded, with flare durations of several days, very high luminosities and very high temperatures (see for example Graffagnino et al. 1995; Preibisch et al. 1995; Güdel et al. 1999). Further, unusual flare shapes have also been reported for some stars (see for examples Haisch et al. 1987; White et al. 1986; Reale et al. 2004; Raassen et
Chapter 1. Introduction

Figure 1.10: Flare observed with Trace using the 171 Å bandpass filter. In the left panel, a bright, flaring loop is seen after about 15 min from the flare onset. In the center, taken 1.5 hours later, much larger loops are visible. In the right panel, the magnetic configuration has simplified and the loops cool down. Several smaller flares accompanied this event.

A clear correlation between the coronal temperature during the flare peak (i.e., the time interval when the largest photon rate is detected, $T_p$) and the peak emission measure ($EM_p$) has been found for both solar and stellar flares. Güdel (2004a) fitted a regression line to the $T_p$ and $EM_p$ values obtained from a sample of stellar flares that had been modeled previously and found a relation $EM_p \propto T_p^{3.30 \pm 0.35}$. A similar relation for solar flares was reported by Feldman et al. (1995), which however does not smoothly connect to the stellar relation (solar flares are found to be lower in $EM$ for a given $T$, or higher in $T$ for a given $EM$). This relation is reminiscent of the temperature-$L_X$ relation found for quiescent main-sequence stars (Güdel 2004a). This analogy would suggest that flares are the major contributors to the X-ray luminosities of stars, so that the more luminous stars are dominated by hotter and larger flares.

Coronal mass ejections (CMEs) are large ($\approx R_\odot$) erupting loop-like features. CMEs can be observed in white light using coronographs. Material from the corona is ejected into interplanetary space. CMEs are associated with flares, but the relation between them is still not fully understood.

1.5.8 X-Ray Emission from Accretion Shocks

The discovery that at least part of the X-ray emission of CTTS could be generated in accretion shocks brings an other important element to our understanding of highly energetic processes in young stars. The first high-resolution spectrum of a CTTS, TW Hya, was presented by Kastner et al. (2002). The Chandra spectrum turned out to be dominated by cool plasma (peaking at $\log T[\text{K}]= 6.5$). Further, density-sensitive He-like triplets of NeIX and OVII indicate very high densities ($\log n_e = 13 \ [\text{cm}^{-3}]$). Abundance anomalies were found, with a deficiency of Fe but an enhanced Ne abundance. Given the soft emission and the high densities, the authors suggested that the bulk of the
X-ray emission in TW Hya is generated in accretion shocks. Stelzer & Schmitt (2004) confirmed this finding from the XMM-Newton high-resolution spectrum of TW Hya. They further interpreted the very high Ne/Fe abundance ratio as a depletion of Fe by condensation into dust grains in the accretion disk. High electron densities were further measured in the high-resolution spectra of other accreting stars like BP Tau (Schmitt et al. 2005; Robrade & Schmitt 2006), and V4046 Sgr (Günther et al. 2006), and were interpreted as indications of X-ray generation in accretion shocks. However, the spectrum of BP Tau is dominated by hot plasma \( (T \approx 20 \text{ MK}) \), that cannot be generated in accretion shocks. Therefore, the hot component is likely to originate from coronal plasma, while the soft component may originate from accretion shocks.

We can estimate the luminosity, the temperature and the density that we expect in the case that the X-rays are generated in accretion shocks. Assuming that the disk is truncated at the corotation radius \( (R_{\text{cor}}) \), the accretion luminosity is

\[
L_{\text{acc}} = (1 - \frac{R}{R_{\text{cor}}}) \frac{GM \dot{M}}{R} \approx 1200 (1 - \frac{\tilde{R}}{R_{\text{cor}}}) \frac{\tilde{M} \tilde{M}_{-8}}{\tilde{R}} \times 10^{30} \text{ erg s}^{-1} \quad (1.10)
\]

where \( \dot{M} = \dot{M}/M_{\odot}, \tilde{M}_{-8} = \dot{M}/10^{-8}M_{\odot}\text{ yr}^{-1}, \tilde{R} = R/R_{\odot}, \) and \( \tilde{R}_{\text{cor}} = R_{\text{cor}}/R_{\odot} \). In case of strong shock \( (n_2 = 4n_1, v_1 = 4v_2, \) where \( n_1 \) and \( v_1 \) are the pre-shock density and velocity, and \( n_2 \) and \( v_2 \) are the post-shock density and velocity), the temperature is given by

\[
T = 3 v_1^2 \mu \frac{m_p}{16k}, \quad (1.11)
\]

where the velocity is approximately the free fall velocity \( v_{\text{ff}} = (2GM/R)^{1/2} \), \( m_p \) is the proton mass, \( k \) is the Boltzmann constant, and \( \mu \) is the mean molecular weight for a fully ionized gas \( (\mu \approx 0.62) \). Further, we can also estimate the shock density. We use the strong shock condition \( n_2 = 4n_1 \). The pre-shock density \( n_1 \) can be estimated from the mass accretion rate and the accreting area on the stellar surface: \( \dot{M} \approx 4\pi R^2 f v_{\text{ff}} n_e m_p, \) where \( f \) is the surface filling factor of the accretion flow, which ranges between 0.1 and 10% according to Calvet & Gullbring (1998). We thus obtain for the post-shock density:

\[
n_2 \approx 4 \times 10^{11} \frac{\tilde{M}_{-8}}{\tilde{R}^{3/2} \tilde{M}_{1/2}^{1/2} f} \text{[cm}^{-3}] \quad (1.12)
\]

Using this formula and measurements of luminosity, temperature, and electron density one can test the hypothesis that X-rays are formed in accretion shocks in CTTS stars. The properties found for TW Hya are consistent with the hypothesis that the bulk of the X-ray emission originates in accretion shocks. For BP Tau, the cool component may also be formed in shocks.
Anomalies in the abundances derived from X-ray spectra are proposed as a further evidence of accretion. As already mentioned above, Stelzer & Schmitt (2004) suggested that the very high Ne/Fe abundance ratios (Ne/Fe = 8-11, referring to the solar photospheric abundances of Anders & Grevesse 1989, and Grevesse & Sauval 1999 for Fe) in the spectrum of TW Hya could be due to metal depletion by condensation onto grains in the accretion disk. (The authors did, however, also not exclude possible environmental abundance anomalies in the original molecular cloud.) In BP Tau, Robrade & Schmitt (2006) measured a Ne/Fe abundance ratio of \( \approx 5 \), which is anomalous, but not as high as in TW Hya. In the TWA 5 system, another member of the TW Hya association, Argiroffi et al. (2005) measured Ne/Fe \( \approx 6 \). Further, Kastner et al. (2004) measured a quite high Ne/Fe abundance ratio in the WTTS HD 98800 (Ne/Fe \( \approx 4 \)). Drake et al. (2005b) proposed to use the Ne/O abundance ratio as a diagnostic for metal depletion in accreting stars. Contrary to the Ne/Fe abundance ratio, the Ne/O ratio is substantially larger in TW Hya than in the other stars in the TW Hya association. However, the Ne/O abundance found for BP Tau is low and similar to the ratios found for non-accreting stars. Drake et al. (2005b) suggested that this is due to evolutionary effects in the circumstellar disk.

In summary, there are three X-ray properties that have been suggested to be signatures of accretion-induced X-ray emission: a soft spectrum, high electron densities, and abundance anomalies (a high Ne/Fe abundance ratio as suggested by Stelzer & Schmitt 2004 or a high Ne/O abundance ratio as suggested by Drake et al. 2005b). A natural question is then if the properties measured in TW Hya and partly in BP Tau are common to all accreting stars.

Recently, Güdel et al. (2007b) measured a low electron density in the O\textsc{vii} triplet of the CTTS T Tau (\( n_e \lesssim 10^{10} \text{ cm}^{-3} \)). The X-ray emitting plasma of the latter star is found to be very hot on the one hand (with coronal temperatures up to \( \approx 30 \text{ MK} \)), but in the XMM-Newton RGS spectrum a large soft excess has also been measured (\( T \approx 1 - 3 \text{ MK} \)). The measured low density cannot be explained in an accretion shock scenario. The spectrum of T Tau is shown in Fig. 1.11 and compared with the spectra of the active HR 1099 RS CVn binary and the inactive F dwarf Procyon. While absorption is negligible in the spectra of HR 1099 and Procyon, it affects the spectrum of T Tau (\( N_H \approx 0.5 \times 10^{22} \text{ cm}^{-2} \), Güdel et al. 2007b), so that the lines at longer wavelengths are strongly attenuated. At wavelengths < 15 Å, the spectrum of T Tau is comparable with the spectrum of the active HD 1099. However, at longer wavelengths, the intrinsic T Tau spectrum is more similar to the spectrum of the inactive Procyon. The O\textsc{vii} triplet is much stronger in the spectrum of T Tau than in HR 1099 and is more similar to the triplet of the Procyon spectrum. The ratio between the O\textsc{vii} triplet and the O\textsc{viii} Ly\( \alpha \) of T Tau is also similar to the ratio observed in Procyon, suggesting a large amount of cool plasma.
1.6. The XMM-Newton Extended Survey of the Taurus Molecular Cloud (XEST)

The new X-ray observatories *Chandra* and *XMM-Newton* provide unprecedented power to study stellar populations embedded in star forming regions. A large survey of the Orion Nebula Cluster was recently performed with *Chandra*. This project, the “*Chandra Orion Ultradeep Project*”, and its results were published in a special issue of the Astrophysical Journal (see the introductory paper by Getman et al. 2005). The survey has monitored a 17′ × 17′ field around the Trapezium in which approximately 1400 young stellar objects are present. Results on the rotation-activity relation and the difference between accreting and non-accreting stars have been reported by Preibisch et al. (2005). In agreement with previous studies of the ONC cloud (Feigelson et al. 2002; Flaccomio et al. 2003a), the authors found that, in contrast to MS stars, magnetic activity does not depend on rotation in the Orion Cluster, rather, all stars lie in the saturation regime. Further, CTTS were found to be 2-3 times less active on average than WTTS and to display larger scatter in correlations of $L_X$ with stellar parameters such as mass or stellar luminosity. While WTTS show properties similar to MS stars, X-ray activity is found to be suppressed.
in CTTS. Different hypotheses for the deficiency of X-ray activity in accreting stars have been proposed, such as a change in the coronal structure, a change of the internal stellar structure, or suppression of plasma heating in mass-loaded magnetic loops.

Three chapters of this thesis (Chapters 3–5, based on Telleschi et al. 2007a,b,c) result from a large survey of the Taurus Molecular Cloud. This project is briefly introduced here.

The Taurus Molecular Cloud (TMC) is the closest and best-studied low-mass star-forming region. The TMC is rather large, covering an area in the sky of 10-15 degrees in diameter, corresponding to 25-35 pc at a distance of 140 pc. In the TMC, star formation occurs in the “isolated mode”, rather than in dense clusters of low-mass stars around massive O-type stars like in the Orion Nebula Cluster. Except for few Herbig Ae/Be stars, the members of TMC are low-mass stars, most of them having masses below one solar mass.

The XMM-Newton Extended Survey of the Taurus Molecular Cloud (XEST) is a large study of the most populated areas of the TMC. About 5 square degrees of the star forming region have been observed in 28 XMM-Newton exposures. The initial XEST project collected 19 fields, each having approximately 30 ks of exposure time. This sample was complemented with observations from other projects obtained from the XMM-Newton archive. These observations usually have longer exposure times (see Güdel et al. 2007a for more details). A map of the TMC region is shown in Fig. 1.12. The grayscale background is an extinction ($A_V$) map (from Dobashi et al. 2005), while the contours show CO emission (from Dame et al. 1987). The small white dots mark the individual TMC members, while the large circles illustrate the XMM-Newton fields of view of the survey (the dashed circles are fields from the additional programs).

The data reduction and source identification processes are described in Güdel et al. (2007a). The detection statistics of the XEST project are given in Table 1.2. Almost all WTTS (98%) and most of the CTTS (85%) have been detected. Most of the undetected stars are strongly absorbed objects or very low-mass stars. More details on the XEST X-ray detection statistics can be found in Güdel et al. (2007a).

Spectral analysis was performed for all detected members using the EPIC spectra. PN spectra were used except in cases for which the PN data were not available, for example if a star fell on a PN CCD gap, or if the PN instrument was not in use. In such cases, the MOS spectra were used. Two different models were fitted to the spectral data in the X-ray spectral analysis software, XSPEC (Arnaud 1996), using the vapec thermal collisional ionization equilibrium model. The first model was composed of one or two plasma components that are in thermal equilibrium (1-$T$ or 2-$T$). The abundances were fixed at values typical for PMS stars or active Zero-Age Main-Sequence (ZAMS) stars (based on results from García-Alvarez et al. 2005; Argiroffi et al. 2004 and results from Chapter 2) and were arranged according to a weak “inverse FIP-
Figure 1.12: Map of the TMC region (from Güdel et al. 2007a). The grayscale background is an extinction ($A_V$) map from Dobashi et al. (2005), while the contours show the CO emission (Dame et al. 1987). The XMM-Newton fields of view of the XEST survey are marked with black circles (dashed circles are exposures from other projects also used for the survey). The white dots are individual TMC members.

effect” (C = 0.45, N = 0.788, O = 0.426, Ne = 0.832, Mg = 0.263, Al = 0.5, Si = 0.309, S = 0.417, Ar = 0.55, Ca = 0.195, Fe = 0.195, Ni= 0.195, with respect to solar photospheric abundances reported by Anders & Grevesse 1989). The spectral components were subject to photoelectric absorption, and the hydrogen column density, $N_H$, was treated as a further fit parameter. The 1-$T$ or 2-$T$ fit methods could introduce biases in highly absorbed spectra, in which
Table 1.2: XEST X-ray detection statistics

<table>
<thead>
<tr>
<th>Object type</th>
<th>Members surveyed</th>
<th>Detections</th>
<th>Detection fraction</th>
</tr>
</thead>
<tbody>
<tr>
<td>0, 1 Protostars</td>
<td>20 (21)</td>
<td>8 (10)</td>
<td>40% (48%)</td>
</tr>
<tr>
<td>2 CTTS</td>
<td>65 (70)</td>
<td>55 (60)</td>
<td>85% (86%)</td>
</tr>
<tr>
<td>3 WTTS</td>
<td>50 (52)</td>
<td>49 (50)</td>
<td>98% (96%)</td>
</tr>
<tr>
<td>4 BDs</td>
<td>16 (17)</td>
<td>8 (9)</td>
<td>50% (53%)</td>
</tr>
<tr>
<td>5 Herbig</td>
<td>2</td>
<td>2</td>
<td>100%</td>
</tr>
<tr>
<td>9 others/unident.</td>
<td>6 (7)</td>
<td>4 (5)</td>
<td>67% (71%)</td>
</tr>
<tr>
<td><strong>Total</strong></td>
<td>159 (169)</td>
<td>126 (136)</td>
<td>79% (80%)</td>
</tr>
</tbody>
</table>

NOTES:
Numbers in parentheses include Chandra observations. A lightly absorbed source near L1551 IRS5 has been discounted from being a true counterpart to this highly embedded protostar (non-detection for XEST-22-040).

the soft component is strongly suppressed or entirely undetected. Therefore, in another approach, the spectra were fitted using a model based on a continuous emission measure distribution (EMD) similar to the EMD found previously for young, active ZAMS stars (see for example Chapter 2). The EMD was approximated by a grid of 20 isothermal components binned to intervals of \( d\log T[K] = 0.1 \), from \( \log T[K] = 6.0 \) to \( \log T[K] = 7.9 \), defining two power laws that intersect at the peak temperature, \( T_0 \). The power law below \( T_0 \) has an index \( \alpha \) that was fixed at \( \alpha = 2 \) (suggested by previous studies, see Figure 1.13), while above \( T_0 \) the index \( \beta \) was free to vary between -3 and +1. A low- and high-temperature cut-off were set to the two power laws, at \( \log T = 6.0 \) and \( \log T = 8.0 \), respectively. Finally, the normalization of our EMD was defined as the emission measure (per 0.1 dex in \( \log T \)) at the peak temperature (EM\( T_0 \)). Abundances were fixed to the same values as in the 2-\( T \) fits. The parameters fitted were thus \( T_0 \), EM\( T_0 \), \( \beta \), and \( N_H \).

X-ray parameters obtained from the spectral fitting and a catalog with fundamental properties of the observed TMC members are given in Tables 4–11 of Güdel et al. (2007a). Results from the survey are presented in a series of papers; here, we briefly list some of the key results.

- A clear correlation has been found between mass and \( L_X \) and between \( L_X \) and the stellar bolometric luminosity \( L_* \) (see Güdel et al. 2007a and Chapter 4). Further, accreting stars are found to be less X-ray active than non-accreting stars, but at the same time CTTS are found to reveal a higher average coronal temperature than WTTS (see Chapter 4).
- The Taurus-Auriga complex is the only star forming region for which
1.6. The XMM-Newton Extended Survey of the Taurus Molecular Cloud (XEST)

Figure 1.13: EMD of the solar analog star 47 Cas B reconstructed from high-resolution X-ray spectroscopy (black histogram; red histograms give upper and lower error ranges). The EMD can be approximated by two power laws with indices $\alpha$ and $\beta$, respectively. The power law below $T_0$ has an index of $\approx 2$. The power laws are displayed as solid lines in the figure, although in our numerical model they are approximated by bins of width $d\log T = 0.1$. $T_0$ is the temperature of the bin at the peak, and $EM_0$ is the emission measure per $d\log T$ in this bin.

an anti-correlation of X-ray activity and rotation period, similar to the relation observed in MS stars, had been previously reported (see, e.g., Stelzer & Neuhauser 2001). Briggs et al. (2007), however, show that this anti-correlation can be explained by on average higher stellar luminosities ($L_*$) and effective temperatures ($T_{\text{eff}}$) of fast rotators together with the almost linear correlation found between $L_X$ and $L_*$.  

- In Chapter 3, results from a sample of nine high-resolution spectra of PMS stars in the Taurus-Auriga complex are presented. A soft excess has been discovered in the high-resolution spectra of the accreting stars while this feature is absent in non-accreting stars.

- X-ray spectra of jet-driving sources, showing two different emission components subject to different photoelectric absorption, are presented by Güdel et al. (2007a). They suggest that dust-depleted accretion streams from the disk to the stars absorb the underlying X-ray coronae.

- Other papers related to the survey include: an optical monitor survey of the TMC (Audard et al. 2007), X-ray studies of brown dwarfs (BDs) (Grosso et al. 2007a), a U-band survey of the BDs (Grosso et al. 2007b), light curve variability studies (Stelzer et al. 2007), spectral analysis of
flares (Franciosini et al. 2007), an investigation of statistical fluctuations in X-ray light curves (Arzner et al. 2007a), an investigation of spectral parameters in faint spectra (Arzner et al. 2007b), an analysis of the gas-to-dust ratio in the TMC (Glauser et al. 2007), a study of the high-resolution spectrum of the Herbig Ae/Be star AB Aur (see Chapter 5), and a search for new TMC members (Scelsi et al. 2007).
Chapter 2

High-Resolution X-Ray Spectroscopy of Solar Analogs with Different Ages

ABSTRACT: We investigate the long-term evolution of X-ray coronae of solar analogs based on high-resolution X-ray spectroscopy and photometry with XMM-Newton. Six nearby main-sequence G stars with ages between \( \approx 0.1 \) Gyr and \( \approx 1.6 \) Gyr and rotation periods between \( \approx 1 \) d and 12.4 d have been observed. We use the X-ray spectra to derive coronal element abundances of C, N, O, Ne, Mg, Si, S, and Fe and the coronal emission measure distribution (EMD). We find that the abundances change from an inverse-First Ionization Potential (FIP) distribution in stars with ages around 0.1 Gyr to a solar-type FIP distribution in stars at ages of 0.3 Gyr and beyond. This transformation is coincident with a steep decline of non-thermal radio emission. The results are in qualitative agreement with a simple model in which the stream of electrons in magnetic fields suppresses diffusion of low-FIP ions from the chromosphere into the corona. The coronal emission measure distributions show shapes characterized by power-laws on each side of the EMD peak. The latter shifts from temperatures of about 10 MK in the most rapidly rotating, young stars to temperatures around 4 MK in the oldest target considered here. The power-law index on the cooler side of the EMD exceeds expected slopes for static loops, with typical values being 1.5–3. We interpret this slope with a model in which the coronal emission is due to a superposition of stochastically occurring flares, with an occurrence rate that is distributed in radiated energy \( E \) as a power-law, \( dN/dE \propto E^{-\alpha} \), as previously found for solar and stellar flares. We obtain the relevant power-law index \( \alpha \) from the slope of the high-temperature tail of the EMD. Our EMDs indicate \( \alpha \approx 2.2 – 2.8 \), in excellent agreement with values previously derived from light curves of magnetically active stars. Modulation with time scales reminiscent of flares is found in the light curves.
of all our targets. Several strong flares are also observed. We use our $\alpha$ values to simulate light curves and compare them with the observed light curves. We thus derive the range of flare energies required to explain the light-curve modulation. More active stars require a larger range of flare energies than less active stars within the framework of this simplistic model. In an overall scenario, we propose that flaring activity plays a larger role in more active stars. In this model, the higher flare rate is responsible both for the higher average coronal temperature and the high coronal X-ray luminosity, two parameters that are indeed found to be correlated.
2.1 Introduction

The solar magnetic field has steadily declined during the Sun’s evolution on the main sequence. Studies of stellar clusters and individual field stars with approximately known ages have shown that the principal parameter determining magnetic activity on a star is its spin rate which, together with convection, controls the operation of the internal magnetic dynamo. As a star spins down due to angular momentum loss via a magnetized stellar wind, its dynamo action weakens, thus continuously reducing magnetic activity expressed in the outer stellar atmosphere. The spin-down history of a solar-like star has been studied in detail by using open-cluster samples (e.g., Bouvier 1990; Soderblom et al. 1993) and accompanying theoretical calculations (e.g., Pinsoneault et al. 1989; MacGregor & Brenner 1991). For a given main-sequence stellar mass in the spectral domain of G and K stars, the rotation period is a rather well-defined function of the star’s age if the latter exceeds a few 100 Myr, regardless of the initial Zero-Age Main-Sequence (ZAMS) rotation rate. For near-ZAMS stars, in contrast, the rotation rate depends on the pre-main sequence evolution and may range between rather modest values (rotation periods $P$ of several days) and rates of the so-called ultra-fast rotators (periods of $\lesssim 1$ d), regardless of the precise stellar age. This spread in rotation rate is well illustrated by stellar cluster samples, such as the Pleiades, in which G and K stars are still close to the ZAMS (Soderblom et al. 1993; Stauffer et al. 1994).

Magnetic activity expresses itself in the outer stellar atmospheres through various phenomena such as dark magnetic spots, bright chromospheric plages, chromospheric emission lines, and coronal radio and X-ray emissions. The coronal emissions display the largest range of variation in response to the surface magnetic activity level. Whereas starspots may cover a few percent of the photosphere of the most active solar analogs, thus producing a photometric wave with a modulation depth of a few percent at best, the range of X-ray luminosity $L_X$ in a solar analog spans at least three orders of magnitude between spun-down inactive examples such as the Sun or $\beta$ Hyi (with $L_X$ between a few times $10^{26}$ erg s$^{-1}$ and a few times $10^{27}$ erg s$^{-1}$) and rapid rotators at the “saturation level” ($L_X \approx 10^{-3} L_{bol} \approx [2 - 4] \times 10^{30}$ erg s$^{-1}$ for a solar-mass star; Maggio et al. 1987; Gudel et al. 1997a). Even the magnetic cycle of the Sun induces an X-ray luminosity variation over one order of magnitude (Micela & Marino 2003). Such systematic cyclic or semi-cyclic variations in turn limit the accuracy with which we can attribute “characteristic” X-ray luminosity levels to stars of a given age if other solar analogs are subject to similar magnetic cycles in X-rays, as seems to be the case (Dorren & Guinan 1994; Dorren et al. 1995).

Short-wavelength ultraviolet and X-ray emissions of a star like the Sun not only serve as a valuable diagnostic to study stellar spin-down and the oper-
ation of the internal dynamo, they are also pivotal for the evolution of the outer atmospheres of planets, in particular the photochemistry in the early atmospheres of Earth-like planets (Canuto et al. 1982; Ribas et al. 2005). Motivated by the interest in understanding the Sun’s and our solar system’s past, we have been studying the “Sun in Time” from the early evolutionary stages on the ZAMS to the terminal stage on the main sequence, at ages of 5-10 Gyr. This study encompasses various wavelength regimes, including radio (Güdel & Gaidos 2001b), optical and ultraviolet (Dorren & Guinan 1994; Guinan et al. 2003), and X-ray wavelengths (Dorren et al. 1995; Güdel et al. 1997a, 1998b). The latter two studies comprise a detailed description of various aspects of the X-ray emission of a solar analog during its main-sequence evolution, based on low-resolution X-ray spectroscopic data from the ROSAT and ASCA satellites. Among the principal findings of that work was a clear trend of decreasing coronal electron temperatures as a solar analog ages, following the decrease in overall X-ray luminosity from ages of \( \approx 0.1 \) Gyr all the way to ages of nearly 10 Gyr. The authors speculated that the decreasing heating efficiency is due to a decreasing level of coronal flaring owing to a smaller filling factor in older stars. In this picture, the flare rate is responsible both for bringing dense material into the corona and heating it to high temperatures, analogous to the behavior of individual flares observed on the Sun. A large flare rate in younger stars with a higher magnetic filling factor could thus produce X-ray coronae that are more luminous and at the same time are hotter. A larger filling factor is, however, not necessarily needed to make the corona more luminous or hotter, as we will discuss in this chapter.

The new generation of X-ray observatories, XMM-Newton and Chandra, offers entirely new access to stellar coronal physics by providing high-resolution X-ray spectroscopy with considerable sensitivity. Apart from more detailed studies of the thermal stratification, they also allow us to derive the coronal abundances of individual elements and to measure electron densities. Both parameters are important for our understanding of the physical processes in stellar coronae. It is now well established that coronae of magnetically active stars show various anomalies in their composition, in particular an overall depletion of metals (e.g., Drake et al. 1994; White et al. 1994) and a relative underabundance of elements with a low (< 10 eV) First Ionization Potential (FIP) compared to elements with a high (> 10 eV) FIP (e.g., Brinkman et al. 2001; Güdel et al. 2001a; Drake et al. 2001). In the inactive Sun, in contrast, element abundances are arranged according to the so-called FIP-effect, with low-FIP elements being overabundant in the corona relative to their photospheric values and relative to the higher-FIP elements (Meyer 1985a,b; Feldman 1992). Similar trends have been noted in inactive stellar coronae (Laming et al. 1996; Drake et al. 1997; Güdel et al. 2002). Because the elemental composition and the element fractionation ultimately derive from the photospheric gas and the physical processes heating it and transporting it to coronal heights,
they maybe important tracers for the coronal heating mechanism. For example, it has been suggested that the Ne enrichment seen in active stars may be related to increased levels of flaring (Brinkman et al. 2001).

This chapter presents a study of six solar analogs at young and intermediate ages observed with the instruments on board XMM-Newton. The stars range from near-ZAMS ages (0.1 Gyr) to ages between approximately 1-2 Gyr when the rotation period has decreased to about 12 days. Together with the Sun, they cover almost the entire path of coronal main-sequence evolution. This study is in many ways a follow-up and continuation of the work presented by Güdel et al. (1997a).

Our work emphasizes the similarity of stellar mass, surface gravities and internal structure, i.e., we confine this study to the main sequence and treat rotation (or equivalently, age) as our principal free parameter. In a complementary study, Scelsi et al. (2005) investigate three G-type stars, all at very high activity levels but at largely differing evolutionary stages and with different internal structure and surface gravities, including the young main-sequence star EK Dra, a pre-main sequence (weak-lined T Tau) star, and the evolved Hertzsprung-gap giant 31 Com. They find very similar emission measure distributions in the latter two active stars, regardless of the differences in their fundamental parameters.

The outline of our presentation is as follows. In Sect. 2.2 and 2.3 we discuss our targets, the observations, and the principal procedures of the data reduction, respectively. Special attention has been paid to the analysis and interpretation of the spectroscopic data in order to recognize the principal potential and the limitations of the spectral inversion, i.e., the derivation of emission measure distributions and the element abundances. In Sect. 2.4 we describe the spectra of the six targets. We discuss our methods in Sect. 2.5. Sect. 2.6 presents our results, while we discuss various features and models in Sect. 2.7. Finally, Sect. 2.8 contains our conclusions.

2.2 Targets

2.2.1 General Properties

XMM-Newton data of six young and intermediate-age solar analog stars have been analyzed. The stars are all of early-to-mid spectral type G on the main sequence. Their ages range from approximately 0.1 Gyr for 47 Cas B and EK Dra to $\approx 1.6$ Gyr for $\beta$ Com. These ages have been determined using various proxies such as rotation periods (for the older targets), or memberships in moving groups of known ages (for the younger targets). We have selected these targets because they have been studied in much detail before (e.g., Dorren & Guinan 1994), have well-measured fundamental parameters,
Table 2.1: Program stars, including a comparison with the Sun.

<table>
<thead>
<tr>
<th>Star</th>
<th>Spec. type</th>
<th>Distance $^a$ (pc)</th>
<th>$P_{rot}$ $^a$ (d)</th>
<th>$\log L_X^b$ (erg s$^{-1}$)</th>
<th>$\log L_X^c$ (erg s$^{-1}$)</th>
<th>$\log L_X^d$ (erg s$^{-1}$)</th>
<th>Age$^e$ (Gyr)</th>
</tr>
</thead>
<tbody>
<tr>
<td>47 Cas B</td>
<td>G0-2 V</td>
<td>33.56</td>
<td>$\approx 1.0$</td>
<td>30.31</td>
<td>30.35</td>
<td>30.39</td>
<td>0.1</td>
</tr>
<tr>
<td>EK Dra</td>
<td>G0 V</td>
<td>33.94</td>
<td>2.75</td>
<td>29.93</td>
<td>30.06</td>
<td>30.08</td>
<td>0.1</td>
</tr>
<tr>
<td>$\pi^1$ UMa</td>
<td>G1 V</td>
<td>14.27</td>
<td>4.7</td>
<td>29.10</td>
<td>29.05</td>
<td>29.06</td>
<td>0.3</td>
</tr>
<tr>
<td>$\chi^1$ Ori</td>
<td>G1 V</td>
<td>8.66</td>
<td>5.1</td>
<td>28.99</td>
<td>28.95</td>
<td>28.95</td>
<td>0.3</td>
</tr>
<tr>
<td>$\kappa^1$ Cet</td>
<td>G5 V</td>
<td>9.16</td>
<td>9.2</td>
<td>28.79</td>
<td>28.94</td>
<td>28.95</td>
<td>0.75</td>
</tr>
<tr>
<td>$\beta$ Com</td>
<td>G0 V</td>
<td>9.15</td>
<td>12.4</td>
<td>28.21</td>
<td>28.26</td>
<td>28.26</td>
<td>1.6</td>
</tr>
<tr>
<td>Sun</td>
<td>G2 V</td>
<td>$5 \times 10^{-6}$</td>
<td>25.4</td>
<td>27.3</td>
<td>27.3</td>
<td>27.3</td>
<td>4.6</td>
</tr>
</tbody>
</table>

$^a$ stellar distances from Perryman et al. (1997).

$^b$determined from ROSAT in the 0.1–2.4 keV band (Güdel et al. 1997a, 1998a,b).

$^c$determined from XMM-Newton in the 0.1–2.4 keV band (this work).

$^d$determined from XMM-Newton in the 0.1–10 keV band (this work).

$^e$from Güdel et al. (1997a, 1998a,b).

and are well-behaved representatives of their age class. A detailed summary and discussion is given in Dorren & Guinan (1994) and Güdel et al. (1997a). The distances quoted are derived from the Hipparcos parallaxes (Perryman et al. 1997), and the rotation periods have been measured photometrically. The general properties of the stars are listed in Table 2.1, where they are also compared with solar values. The X-ray luminosities $L_X$ given there refer to measurements from ROSAT photometry in the 0.1–2.4 keV band (Güdel et al. 1997a, 1998a), and to the spectral modeling discussed in the present work, also for the 0.1–2.4 keV band, and additionally for the 0.1–10 keV band. For the X-ray luminosity of the Sun, we use a representative value of $\log L_X = 27.3$ as in Güdel et al. (1997a). This value is in agreement with the $L_X$ of α Cen determined by Raassen et al. (2003), $L_X = 1.6 \times 10^{27}$ erg s$^{-1}$. A steady decline of $L_X$ with increasing age and increasing rotation period $P$ is evident.

The most active target, 47 Cas B, is the fainter component in the 47 Cas binary system. It has not been characterized optically as it has been individually detected only by radio methods (Güdel et al. 1998b) and indirectly from Hipparcos measurements as a companion to the optically bright F0 V star 47 Cas = HR 581 = HD 12230. The radio position is clearly offset from the position of the F0 star. It is, furthermore, very unlikely that early F stars display luminous and spectrally hard X-ray emission (Panzera et al. 1999). On the other hand, all characteristics of the X-ray detection fit well to an early G-type star with an age similar to that of the Pleiades. In particular, the X-ray luminosity corresponds to the saturation level of an early G star ($L_X \approx 10^{-3} L_{bol}$, Vilhu 1984), thus excluding a corona of a later-type star as its origin.

A saturated corona requires, for a G star, a rotation period smaller than
that of EK Dra. A periodic signal with a period of $\approx 1$ day was reported from the ROSAT All-Sky Survey observations, and was attributed to stellar rotation (Güdel et al. 1995). This period also coincides with the rotation period of the fastest early-G type rotators in the Pleiades (Soderblom et al. 1993). We refer the reader to the detailed discussion in Güdel et al. (1995) and Güdel et al. (1998b). We tried to quantify the maximum contribution of the F0 star to the X-ray luminosity of the 47 Cas system. ROSAT studies of the Pleiades (Stauffer et al. 1994; Micela et al. 1996, 1999) found only few stars with spectral type A7-F3 that have $\log L_X > 29.4$ (i.e., more than 10% of the $L_X$ of 47 Cas), and all of them have known late-type companions, which are likely to be the sources of the high X-ray luminosity (Mermilliod et al. 1992). In addition, the typical X-ray spectrum of F-type stars in the Pleiades is softer than that of G-type stars (Gagne et al. 1995), so we expect the F primary to provide even less flux from hot plasma. In summary, we thus expect the F0 star to contribute less than 10% to the total detected X-ray flux of the 47 Cas system. We therefore add this target to our list of G-type stars, being the only likely solar analog that is accessible to high-resolution X-ray spectroscopy and that is more active than EK Dra, in fact reaching the saturation level.

### 2.2.2 Photospheric Composition

When measuring coronal abundances, $A$, of elements, a principal problem is to what standard set of abundances the results should refer to. Often, stellar coronal abundances are cited with respect to the solar photospheric composition. But because the stellar coronal plasma ultimately originates in the respective stellar photosphere, coronal abundances may reflect the composition of the latter, and abundance “anomalies” may simply be due to a non-solar composition of the stellar photospheric material. Fortunately, the photospheric abundances of most of our objects have comprehensively been measured. The reported photospheric abundances of our sample of stars (except 47 Cas) are given in Table 2.2. The entries refer to individual measurements or to catalogs compiled from previous studies.
Table 2.2: Stellar photospheric abundances. Values refer to the solar photospheric composition. If available, error ranges are given in parentheses.

<table>
<thead>
<tr>
<th>Star</th>
<th>Fe</th>
<th>Mg</th>
<th>Si</th>
<th>S</th>
<th>C</th>
<th>O</th>
<th>N</th>
<th>Ref.</th>
</tr>
</thead>
<tbody>
<tr>
<td>EK Dra</td>
<td>1.20</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>1</td>
</tr>
<tr>
<td>π¹ UMa</td>
<td>0.83-1.02</td>
<td>0.65-0.83</td>
<td>0.78-1.12</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>2</td>
</tr>
<tr>
<td></td>
<td>0.93 (0.81-1.07)</td>
<td>0.74 (0.58-0.95)</td>
<td>0.89 (0.83-0.95)</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>3</td>
</tr>
<tr>
<td></td>
<td>0.87</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>1</td>
</tr>
<tr>
<td></td>
<td>1.10</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>4</td>
</tr>
<tr>
<td></td>
<td>0.83-0.98</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>5</td>
</tr>
<tr>
<td></td>
<td>1.09 (1.00-1.19)</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>6</td>
</tr>
<tr>
<td>χ² Ori</td>
<td>0.89-0.93</td>
<td>0.91</td>
<td>0.98</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>7</td>
</tr>
<tr>
<td>0.91</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>8</td>
</tr>
<tr>
<td>≈ 1</td>
<td>-</td>
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<td>-</td>
<td>0.63</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>9</td>
</tr>
<tr>
<td>1.14 (1.07-1.22)</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
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<td>-</td>
<td>-</td>
<td>6</td>
</tr>
<tr>
<td>1.35</td>
<td>-</td>
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<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>1</td>
</tr>
<tr>
<td>0.66-1.29</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>5</td>
</tr>
<tr>
<td>κ Or</td>
<td>0.89-1.04</td>
<td>0.85-0.98</td>
<td>0.95-1.07</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>2</td>
</tr>
<tr>
<td>1.29 (1.17-1.41)</td>
<td>0.91 (0.76-1.10)</td>
<td>0.85 (0.76-0.93)</td>
<td>-</td>
<td>0.91 (0.71-1.17)</td>
<td>-</td>
<td>-</td>
<td>3</td>
<td></td>
</tr>
<tr>
<td>1.0</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>7</td>
</tr>
<tr>
<td>1.13</td>
<td>-</td>
<td>-</td>
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<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>6</td>
</tr>
<tr>
<td>1.66</td>
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<td>1</td>
</tr>
<tr>
<td>0.98-1.10</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>5</td>
</tr>
<tr>
<td>β Com</td>
<td>0.93</td>
<td>-</td>
<td>-</td>
<td>1.38</td>
<td>0.98</td>
<td>1.26 (1.05-1.47)</td>
<td>1.05</td>
<td>10</td>
</tr>
<tr>
<td>1.00-1.07</td>
<td>1.17</td>
<td>1.00</td>
<td>-</td>
<td>-</td>
<td>-</td>
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<td>7</td>
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<td>1.15 (1.07-1.25)</td>
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<td>-</td>
<td>-</td>
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<td>-</td>
<td>-</td>
<td>6</td>
</tr>
<tr>
<td>1.07</td>
<td>-</td>
<td>-</td>
<td>-</td>
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<td>-</td>
<td>-</td>
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<td>1</td>
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<tr>
<td>1.00</td>
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<td>-</td>
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<td>-</td>
<td>8</td>
</tr>
<tr>
<td>1.17</td>
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<td>-</td>
<td>-</td>
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<td>4</td>
</tr>
<tr>
<td>0.89-1.17</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>-</td>
<td>5</td>
</tr>
</tbody>
</table>

(1) Rocha-Pinto et al. (2004); (2) Ottmann et al. (1998); the original values were transformed to solar abundances as given by Anders & Grevesse 1989 = AG89 except for Fe for which we use the value given in Grevesse & Sauval 1999 = GS99; (3) Gaidos & Gonzalez (2002); (4) Gray et al. (2001); (5) Cayrel de Strobel et al. (2001); (6) Taylor (2003), corrected to GS99; (7) Edvardsson et al. (1993); (8) Gratton et al. (1996); (9) Tomkin et al. (1995), corrected to AG89 and GS99; (10) Clegg et al. (1981).
2.3. Observations and Data Reduction

As one sees from this literature survey, almost all Fe abundances reported for our targets are compatible with solar photospheric values. Moreover, several measurements for other elements exist, including high-FIP elements such as C, N, and O, and again most of the reports are compatible with solar abundances. We find a trend for super-solar abundances in $\beta$ Com, by perhaps 10–20%. Nevertheless, the photospheric-abundance tabulations for our targets are incomplete and the given measurements scatter, making it impossible to reliably express our coronal abundances relative to the respective photospheric values. However, the above summary makes it clear that our targets (except for 47 Cas) must show a near-solar composition, and we therefore adopt the solar photospheric abundances as the reference composition throughout the chapter. As for 47 Cas, to our knowledge no photospheric abundances of the brighter F star in the system have been reported. However, this system is young and is a member of the Local Association (Güdel et al. 1998b), for which we can reasonably assume near-solar abundances, similar to EK Dra, which is also a member of the Local Association.

2.3 Observations and Data Reduction

2.3.1 Observations

Our target stars were observed with the Reflection Grating Spectrometers (RGS, den Herder et al. 2001) and the European Photon Imaging Cameras (EPIC, Strüder et al. 2001; Turner et al. 2001) on board XMM-Newton (Jansen et al. 2001). The RGS are suited for high-resolution spectroscopy in the wavelength range between 6–38 Å, with a resolution of $\Delta \lambda \approx 60 – 76$ mÅ, hence a resolving power of $\lambda/\Delta \lambda \approx 100 – 500$. The EPICs observe between $\approx 0.15$–$15$ keV, providing a moderate energy resolution of approximately $E/\Delta E = 20 – 50$.

We provide a detailed log of the observations in Table 2.3.

2.3.2 Data Reduction

The data were reduced using the Science Analysis System (SAS) version 5.4.1. We applied the standard processing performed by the RGS metatask rgsproc and the EPIC MOS task emchain. The PN data (used only for the light curves) were reduced using the PN task epchain.

For the RGS data reduction, we extracted the first-order spectra from a spatial cut including 95% of the cross-dispersion Point Spread Function (PSF) (xpsfinc1 = 95 in rgsproc) and an energy cut including 95% of the pulse-height distribution (pdistincl = 95). The background spectra were extracted above and below the source spectra, by excluding 97% of the source
Table 2.3: Observation log

<table>
<thead>
<tr>
<th>Star</th>
<th>Instruments</th>
<th>Filter</th>
<th>Start</th>
<th>Stop</th>
<th>Exposure [s] (^a)</th>
</tr>
</thead>
<tbody>
<tr>
<td>47 Cas B</td>
<td>RGS 1</td>
<td>-</td>
<td>2001/09/10 23:30:24</td>
<td>2001/09/11 13:39:00</td>
<td>36610</td>
</tr>
<tr>
<td>EK Dra</td>
<td>RGS 1</td>
<td>-</td>
<td>2000/12/30 14:01:58</td>
<td>2000/12/31 05:17:04</td>
<td>41960</td>
</tr>
<tr>
<td></td>
<td>RGS 2</td>
<td>-</td>
<td>2000/12/30 14:01:58</td>
<td>2000/12/31 05:17:04</td>
<td>41960</td>
</tr>
<tr>
<td></td>
<td>MOS 2 Thick</td>
<td>-</td>
<td>2000/12/30 14:10:36</td>
<td>2000/12/31 04:38:02</td>
<td>41960</td>
</tr>
<tr>
<td>(\chi^1) Ori</td>
<td>RGS 1</td>
<td>-</td>
<td>2001/04/07 08:56:49</td>
<td>2001/04/07 22:31:53</td>
<td>29326</td>
</tr>
<tr>
<td></td>
<td>MOS 2 Thick</td>
<td>-</td>
<td>2001/04/07 09:03:11</td>
<td>2001/04/07 17:45:10</td>
<td>29326</td>
</tr>
<tr>
<td>(\kappa^1) Cet</td>
<td>RGS 1</td>
<td>-</td>
<td>2002/02/09 16:13:01</td>
<td>2002/02/10 03:21:41</td>
<td>35920</td>
</tr>
<tr>
<td></td>
<td>RGS 2</td>
<td>-</td>
<td>2002/02/09 16:13:01</td>
<td>2002/02/10 03:21:35</td>
<td>35920</td>
</tr>
<tr>
<td></td>
<td>MOS 1 Thick</td>
<td>-</td>
<td>2002/02/09 16:19:33</td>
<td>2002/02/10 03:16:09</td>
<td>35920</td>
</tr>
<tr>
<td></td>
<td>MOS 2 Thick</td>
<td>-</td>
<td>2003/07/20 02:09:06</td>
<td>2003/07/20 19:12:36</td>
<td>61320</td>
</tr>
</tbody>
</table>

\(^a\) Exposure time used for the analysis, in seconds (excluding flares).

For all stars except \(\pi^1\) UMa, \(\chi^1\) Ori, and \(\beta\) Com the MOS data were taken in the small-window mode. In this mode, only the central 100 \(\times\) 100 pixels of the middle CCD are operational, whereas the outer CCDs work in full mode. The small-window mode allows for shorter integration times for the source and thus avoids the brighter sources from becoming piled up. The disadvantage of this mode is, however, that the central window is too small to reliably extract a background region. We therefore selected a source-free region in one of the outer CCDs. The target source itself was extracted from a circle as large as possible in the small window (with a radius of approximately 50\(\arcsec\)).

For \(\pi^1\) UMa, MOS1 was in full window mode whereas MOS2 was in small-window mode. The exposure time was 5000 s longer in MOS1. Pile-up was studied for both MOS cameras with the SAS `epatplot` task. This task makes use of the relative ratio of single- and double-pixel events, that deviate in case of pile-up. Since no evidence for pile-up was found for this star, we decided to use MOS1.

For \(\chi^1\) Ori, MOS1 data were taken in timing mode and MOS2 data in full-window mode. We did not use data taken in timing mode because reliable background subtraction cannot be performed. We found that pile-up was not negligible in the MOS2 data and therefore used an annular extraction region for the source, with inner and outer radius of 80 and 1200 detector pixels (4 and 60 arcsec), respectively, in order to remove the piled-up central part of the PSF.

For \(\beta\) Com both MOS1 and MOS2 observed in full window mode. We
found that pile-up did not affect this source.

In those cases where the background area was selected in one of the outer CCDs, the background count rates could have been underestimated with respect to the source extraction region due to vignetting. We studied background behavior effects using the `eboxdetect` SAS task. This task detects all sources in the field of view; these were subsequently cut out from the image. We then quantified the remaining background count rate per unit area as a function of the distance from the center of the field of view, by analyzing the average rates in annuli at different radii. We found a decrease of at most 10% from the image center to the central portions of the outer CCDs, which is negligible when compared to calibration uncertainties, particularly in the light of the high count rates of our sources. The only possible but small influence could be on the data at the highest energies where the source flux drops below the background flux. We therefore decided not to consider spectral data at energies where the background count rate exceeded the source count rate.

We used the `rgsrmfgen` task for RGS and the `rmfgen` and `arfgen` tasks for MOS to generate response and ancillary files appropriate for the specified extraction regions. To avoid bins with very few counts, we grouped the spectral channels: the MOS data were grouped to a minimum of 25 counts per bin, whereas the RGS data were grouped to a minimum of 10 counts per bin. The very weak continuum level in the RGS spectra of \( \beta \) Com made this grouping scheme not well suited; instead, we regularly rebinned the entire RGS spectra by a factor of five.

Finally, we studied the agreement between the observed wavelengths of bright emission lines with their tabulated rest wavelengths (stellar radial velocities being negligible); in cases where small systematic offsets due to residual calibration uncertainties of order 10 mÅ were found, we slightly updated the assumed boresight coordinates accordingly and repeated the data reduction.

### 2.4 Description of the Spectra

Before quantitatively evaluating the observed spectra, we discuss some general features. Fig. 2.1 reveals four significantly different types of spectra: i) the most active star, 47 Cas B, shows typical features of a very hot plasma, namely a well-developed bremsstrahlung continuum, lines of Mg X and XII, and comparatively high flux ratios of O VIII \( \lambda 18.97/\lambda 21.60 \), of Ne X \( \lambda 12.13/\lambda 13.45 \), of Fe XVIII \( \lambda 14.20/\lambda 15.01 \). ii) The spectrum of EK Dra is significantly cooler, which is in particular evident from the smaller O VIII/O VII flux ratio, a more modest continuum level compared to the line strengths, and a marked dominance of the Fe XVII lines. iii) The third group consists of the intermediately active, intermediately old stars \( \pi^1 \) UMa, \( \chi^1 \) Ori, and \( \kappa^1 \) Cet which all show very similar spectra in which the Fe XVII lines are much stronger than
Figure 2.1: Fluxed, coadded RGS 1 & 2 spectra of the six solar analogs, ordered from high (top) to low (bottom) activity. Examples of error bars at the wavelength of Fe$^{xvii}$ and at $\lambda=20$ Å, a nearly line-free region, are overplotted.
2.4. Description of the Spectra

Figure 2.2: MOS spectra of the six solar analogs, normalized to a distance of 33.56 pc, the distance of 47 Cas B. The overall luminosity of the corona decreases from top to bottom as indicated by the labels.

...those of Fe XVIII, and the O VIII/O VII flux ratio is further reduced. Note that the maximum formation temperature of Fe XVII is only \( \approx 5 \) MK. The continuum has become very weak in these spectra. iv) Finally, the spectrum of the least active target, \( \beta \) Com, is largely dominated by lines of Fe XVII, and the O VIII/O VII flux ratio approaches unity.

Turning to the EPIC MOS spectra in Fig. 2.2, further indicators support this picture. In this figure, we have renormalized the MOS spectra so as to represent the stars at a common distance of 33.56 pc, identical to the distance of 47 Cas B. A marked decrease of the overall emission level is seen as the stellar age proceeds from the sample of Pleiades-age stars (47 Cas B, EK Dra) to the older sample. The most active stars show shallower continuum spectra...
between $\approx 2$ and 10 keV than the less active targets, indicative of the higher overall plasma temperatures of the former. The dominant Fe\textsuperscript{XVII} lines (at 0.826 keV and 0.727 keV) are well developed in the spectra of the less active stars.

### 2.5 Data Analysis

Because the novel aspect of our X-ray data is the high spectral resolution of the RGS, allowing us to access individual emission lines from different elements, we used the spectral data from both RGS instruments but used EPIC data only in so far as they contribute additional information not accessible by RGS, such as spectral data at wavelengths shorter than 6 Å. To keep maximum weight in our data analysis on the RGS data, we restricted the more sensitive EPIC spectral data to one of the three cameras. We chose MOS1 or MOS2, because the MOS spectral resolution is superior to that of the pn camera, and the S/N at higher energies balances well with the S/N provided by RGS at lower energies. Also, the cross-calibration between MOS and RGS is better understood (they use the same mirrors). Only for the light curves did we make use of the data of other EPIC cameras as well.

For the analysis of each target, we used the exposure time covered simultaneously by all three detectors. Each light curve except that of $\beta$ Com contained one well-developed flare. In order to avoid systematic bias by any of these flares (i.e., increasing the average $L_X$ and possibly also increasing the characteristic coronal temperatures), we eliminated the flare intervals from consideration. We will, however, briefly present and discuss the individual light curves separately in Sect. 2.6.3. The exposure times remaining for our spectral analysis are listed in Table 2.3 (last column).

### 2.5.1 Spectral Inversion

An observed coronal X-ray spectrum is the superposition of the spectra emitted by various coronal features with different temperatures, volumes, densities, and possibly different composition in terms of chemical elements. The inversion of an observed spectrum to obtain the underlying physical parameters is therefore a highly degenerate problem, with numerous solutions describing essentially the same spectrum.

For our analysis we consider a coronal model with the following, observationally tested assumptions of the physical parameters. First, the plasma is assumed to be in collisional ionization equilibrium, a model that appears to be sufficiently good as long as no very rapid change in the heating rate of the plasma is taking place (Mewe 1999). We also assume that the coronal plasma is effectively optically thin and that lines with high oscillator strengths
are not subject to resonance scattering, an assumption that has been shown to be justified for the stars in our sample (Ness et al. 2003b). Furthermore, the density-dependence of the populations of metastable levels is neglected, i.e., the spectrum is computed in the low-density limit. This approximation is supported by the flux ratios in the He-like triplet of O\textsuperscript{vii} in our observations, and it appears to be a reasonable assumption for most coronal plasmas (Ness et al. 2004; Testa et al. 2004).

At this point, then, the observed spectrum is essentially a function of the distribution of optically thin coronal features in volume, temperature, and elemental composition; conventionally, the thermal distribution is described by the differential emission measure distribution (DEM),

\[ Q(T) = \frac{n_e n_H dV}{d\ln T} \]  

(2.1)

where \( n_e \) and \( n_H \) are the electron and hydrogen number densities, respectively, and \( dV \) is a differential volume element at temperature \( T \). The DEM determines the line flux \( F_j \) of any given emission line \( j \) through

\[ F_j = \frac{1}{4\pi d^2} \int A(\varphi_j(T))Q(T)d\ln T. \]  

(2.2)

Here, \( \varphi_j \) is the line power per unit emission measure (“emissivity” henceforth), and \( A \) is the abundance of the element emitting the relevant line, with respect to some standard tabulation (such as “solar photospheric abundances”) used for the computation of \( \varphi_j \).

Non-solar abundances in stellar coronae are now well established (e.g., Brinkman et al. 2001; Drake et al. 2001; Audard et al. 2003a), and we recall that the corona of the Sun itself shows considerable deviations from the photospheric mixture. However, it is also known that the solar coronal abundances vary greatly from feature to feature, some showing a marked FIP bias, with others showing photospheric composition (Feldman 1992; Laming et al. 1995). Our spectra are - like most existing stellar X-ray spectra - insufficient to characterize abundances at various separate temperatures; therefore, we will assume \( T \)-independent abundances to at least recognize overall trends in our stellar sample.

The spectral inversion problem is mathematically ill-posed. Statistical scatter due to photon statistics in line fluxes, even if amounting to only a few percent, may introduce considerable scatter in the reconstructed emission measure distribution (Craig & Brown 1976). We expect a strong amplification of such effects from the uncertainties in the atomic physics parameters (which may amount up to a few tens of percent for the emissivities) and from calibration uncertainties (up to several percent in certain spectral regions).

Therefore, the spectral inversion can essentially be performed in a meaningful way only by constraining the problem suitably, for example by subjecting
the reconstruction to smoothness conditions for the resulting, discrete Emission Measure Distribution (EMD), or by iterating the problem using a pre-defined convergence algorithm that reconstructs a preferred, physically meaningful solution (see, e.g., Kaastra et al. 1996a and Güdel et al. 1997b for a discussion of various algorithms with different bias).

We have chosen here to perform the inversion using two widely different algorithms that we briefly summarize as follows:

- **Method 1 (M1).** We fitted the spectrum using synthetic template spectra calculated for a set of physical parameters; the parameters were varied until the fit was optimized. This method has conventionally been used for low-resolution spectra but also for high-resolution grating spectra, in particular as implemented in the SPEX (Kaastra et al. 1996b) and the XSPEC (Arnaud 1996) software packages. We, however, modified the conventional approach by including almost only segments of the spectrum that are dominated by bright lines for which comparatively robust atomic physical parameters should be available, together with some nearly line free regions.

- **Method 2 (M2).** Here, we worked with a list of discrete line fluxes that we extracted from the observed spectrum. If the formation temperatures of these lines cover the range of relevant coronal electron temperatures, then the EMD can be reconstructed by successive approximation, essentially inverting a system of equations like Eq. (2.2).

We emphasize the complementarity of our two approaches: method 1 uses all tabulated emission lines and their blends within the selected spectral segments, while we confine method 2 to a minimum number of lines, namely the brightest lines available, required to derive meaningful EMDs and the most essential abundances.

Both methods are subject to separate biases. Some of the many lines considered in method 1 may be poorly fitted, owing to discrepancy in the tabulated emissivities. If the latter scatter around their true values, the resulting fit may show reduced systematic errors compared to a fit based on one single line. The individual lines in method 2 may be better described, but their small number makes the inversion process rather sensitive to any systematic uncertainty in any of the line emissivities, and some a priori estimate for blend contribution must be considered. Lines not tabulated in the line-emissivity lists affect both methods, either by altering the measured line flux in method 2 through unrecognized blending, or by adding a continuum-like base level if many weak lines contribute in either of the analysis methods.

A comparison of method 2 with a method using polynomial EMDs to calculate synthetic spectra has previously been presented by Audard et al. (2004) for the extremely active and hot corona of the FK Com-type star YY Men.
These authors found that both the reconstructed EMDs and the derived abundances closely agreed for the two methods. Here, we study and compare the results for stars across a large range of activity and coronal temperatures.

To further assess potential sources of systematic error, we applied two different databases to each of the two methods. The spectral models are based on the assumption of coronal ionization equilibrium (CIE). We used the atomic parameters from the MEKAL database in SPEX vers.2.0 (Mewe et al. 1985; Kaastra et al. 1996b) and from the Astrophysical Plasma Emission Code vers.1.3.1 (APEC, Smith et al. 2001) in XSPEC (Arnaud 1996). The MEKAL emissivities are computed using the ionization balance of Arnaud & Raymond (1992) for Fe and Arnaud & Rothenflug (1985) for the other elements. The APED emissivities (emissivities used in the APEC code) are computed using the ionization balance of Mazzotta et al. (1998).

### 2.5.2 Method 1: Synthetic Spectra

Here, we used an approach essentially identical to the one described in Audard et al. (2003a), apart from slightly different wavelength ranges (see below). Because numerous emission lines are poorly described by the spectral line lists in use, we confined our analysis to a restricted number of spectral regions that contain bright lines for which the atomic physics is believed to be relatively well known, and some nearly line-free regions of the spectrum where the continuum dominates. The wavelength ranges of the selected regions are summarized in Table 2.4. As we decided to put more weight on the data from the RGS with its higher spectral resolution, we used EPIC MOS data only in the wavelength range shortward of 9.35 Å, where the RGS effective area is small and the calibration is less accurate. We thus cut the RGS data shortward of 8.3 Å, so that the data from the RGS and the EPIC instruments overlapped around the Mg xi and Mg xii lines. As mentioned in Section 4, we discarded the high-energy part of the EPIC spectrum where the background flux exceeded the source flux (see Table 2.4).

The physical model was defined as a combination of 10 optically thin, thermal CIE models, and a photoelectric absorption component. The photoelectric absorption was frozen by defining a fixed interstellar hydrogen column density. The latter ranged between negligible values (< 10^{18} cm^{-2}) for the closer stars and 7 × 10^{18} cm^{-2} for 47 Cas B. These values are consistent with hydrogen column densities given by Güdel et al. (1997b) and Audard et al. (1999) and with values of the interstellar hydrogen density given by Paresce (1984). However, even the largest values used here were too small to significantly alter our model spectra even at the longest wavelengths in the RGS.

Each of the ten CIE temperatures was confined to within one of ten adjacent temperature intervals of equal width (0.22 dex) in log T; to optimize the multi-temperature fit, given the relatively broad bins, we treat the temperature
Table 2.4: Spectral wavelength ranges used for method 1.

<table>
<thead>
<tr>
<th>Instrument</th>
<th>$\lambda$ range (Å)</th>
</tr>
</thead>
<tbody>
<tr>
<td>RGS 8</td>
<td>30.9 - 50.0</td>
</tr>
<tr>
<td>RGS 12</td>
<td>00.13 - 95.95</td>
</tr>
<tr>
<td>RGS 14</td>
<td>15.15 - 90.0</td>
</tr>
<tr>
<td>RGS 16</td>
<td>20.17 - 15.0</td>
</tr>
<tr>
<td>RGS 17</td>
<td>80.18 - 30.0</td>
</tr>
<tr>
<td>RGS 18</td>
<td>75.19 - 20.0</td>
</tr>
<tr>
<td>RGS 20</td>
<td>80.21 - 10.0</td>
</tr>
<tr>
<td>RGS 21</td>
<td>40.22 - 40.0</td>
</tr>
<tr>
<td>RGS 23</td>
<td>65.24 - 00.0</td>
</tr>
<tr>
<td>RGS 24</td>
<td>50.24.90</td>
</tr>
<tr>
<td>RGS 28</td>
<td>50.30 - 10.0</td>
</tr>
<tr>
<td>RGS 31</td>
<td>10.32 - 00.0</td>
</tr>
<tr>
<td>RGS 33</td>
<td>40.33 - 85.0</td>
</tr>
<tr>
<td>MOS 1</td>
<td>70.6 - 90.0</td>
</tr>
<tr>
<td>MOS 7</td>
<td>80.9 - 35.0</td>
</tr>
</tbody>
</table>

*For 47 Cas B and EK Dra. For the other stars, we used lower limits as follows: 4.96 Å for $\pi^1$ UMa and $\chi^1$ Ori, 4.13 Å for $\kappa^1$ Cet, and 5.0 Å for $\beta$ Com.*

values as free parameters within the bounds of the respective bin intervals, as well as the associated emission measures. The temperature intervals covered the range from $\log T=6$ to $\log T=8.2$ ($T$ is given in K). The abundances of C, N, O, Ne, Mg, Ar, Si, and S were also used as free parameters. Only the abundances found with the 10-temperature fit were used in a second step for the DEM reconstruction. The Chebychev polynomial DEM code in SPEX, with polynomials of order 6 and 8, was used to describe the DEM that best fitted the given spectrum. An example of a best fit is shown in Fig. 2.3, which also illustrates the spectral ranges used for this procedure.

### 2.5.3 Method 2: Inversion of Line-Flux Lists

**EMD Reconstruction**

We reconstructed the discrete EMD starting from fluxes of individual lines. In order to obtain an EMD independently of the element abundances, we selected a few well-defined, bright Fe lines from the ionization stages of Fe$^{XVII}$ to Fe$^{XXV}$ (as far as measurable, see Table 2.5). The Fe$^{XVII-XXIV}$ lines were extracted from the RGS spectrum, whereas the unresolved Fe$^{XXV}$ line triplet was extracted as a single blend system from the MOS spectrum. The emissivities of the Fe$^{XVII-XXIV}$ lines cover the temperature range between 3 MK
2.5. Data Analysis

Figure 2.3: **Upper panel:** data and fitted spectrum for 47 Cas B using method 1 with APEC; **lower panel:** same for method 2. The synthesized spectra from the best-fit parameters are shown in black for the ranges that were used in method 1. Error bars are displayed only for the MOS data in order to avoid confusion. The insets show the important spectral portion containing the Fe L shell lines. The lower plots in each figure show the contributions of each bin to the $\chi^2$ value. Note that for method 2, we show the identical layout for illustration and comparison purposes although the iteration procedure did not make use of a binned spectrum.

and 20 MK rather well, and for the more active stars (47 Cas B and EK Dra), Fe XXV constrains the hottest temperatures, up to about 100 MK (Figure 2.4). In order to obtain information on the cooler part of the EMD in which no Fe lines are detected, we used the flux ratio between the O viii $\lambda$18.97 and the O vii $\lambda$21.60 resonance lines, which itself is independent of abundances. The lines used for each star are listed in Table 2.5 and Table 2.6, together with their measured luminosities.

To avoid cross-talk with the abundance determination, we selected only lines of Fe that are not strongly blended with lines of other elements (with the exception of Fe xix, see below; the Fe XVII lines at 17 Å were not used because they are partly obscured by one of the RGS CCD gaps). Nevertheless, our line
features usually still contain a number of weaker blends from various Fe ions. We thus considered all Fe blending lines close to the principal line as being part of our Fe feature and we therefore computed new $T$-dependent emissivity curves for each such Fe blend system. The line feature, composed of all Fe lines in the chosen wavelength range, was thus subsequently treated like a single emission line. In the RGS spectrum, all tabulated blending Fe lines within ±0.06 Å (one FWHM on each side) of the wavelength of the principal line were considered. In the MOS spectrum, the range was chosen to be ±0.15 Å.

In a first step, we extracted the fluxes of the relevant Fe lines and of the O\textsc{viii} $\lambda$18.97 and O\textsc{vii} $\lambda$21.60 resonance lines from the spectrum. This can be problematic, because the lines have broad wings due to the RGS PSF, and the determination of the underlying continuum is difficult. Moreover, in some cases, blends with lines from different elements may still be present, in particular in the case of the brighter lines of Fe\textsc{xix} that are blended with Ne\textsc{ix} lines. While the effect of these blends cannot be assessed a priori without knowledge of the thermal structure and the abundances, we can obtain useful approximations as follows. Because most ions are represented by several different emission lines, and the continuum can be interpolated from values obtained in nearly line-free regions of the spectrum, a good approximation of the contaminating blends and of the baseline continuum level in fact comes from the fit we derived in method 1. We emphasize that we have not used that model to derive parameters, but only to obtain an approximate description of the baseline spectral flux distribution on which the line fluxes of interest are
superimposed. The latter being the fluxes of Fe lines, we set the Fe abundances in the best-fit model from method 1 to zero, retaining only the emission lines and the continuum from all other elements. We note that neglecting Fe does slightly change the continuum as well, but the influence is of order a few percent in the temperature and wavelength range of interest if solar abundances are assumed. Anticipating sub-solar abundances in active stars with a strong continuum as reported in the previous literature, the contribution of Fe to the continuum will be even smaller. The errors thus introduced into the line-flux extraction are negligible. An equivalent procedure was then also applied to extract the O\textsc{viii} and O\textsc{vii} line fluxes, although in these cases, blending was not significant.

Before extracting individual lines, we need to make sure that the baseline continuum level has been correctly fitted individually in each spectrum. We checked the relevant fits in a nearly line-free region of the RGS spectrum, namely immediately shortward and longward of the O\textsc{viii} $\lambda 18.97$ resonance line by multiplying the EM by a suitable factor. We found that an optimization of the continuum level in these wavelength regions required a change in flux of only a few percent. It is not clear whether this is due to slight cross-calibration problems or to some bias in the fit. Such bias may be introduced by the fact that the MOS spectral portion, determining mostly the hotter part of the DEM, has a relatively high S/N ratio which may lead to some over- or underestimation of the hot EM. Because the latter produces continuum at all wavelengths, a slight offset could also affect the soft part of the RGS spectrum. Given the small magnitude of the effect, we are unable to ascribe such an offset either to (possibly wavelength-dependent) cross-calibration problems or to fit bias as described above. Our final results indicate that any such effect is of minor consequence and does not need to be considered for the overall fit in method 1. For the extraction of line fluxes in method 2, however, we need to optimize the continuum level so that the line flux can be properly defined as an excess above this baseline level.

In a similar manner, we extracted the Fe\textsc{xxv} $\lambda 1.85$ line by first optimizing the nearby continuum level starting from the overall best-fit provided by method 1. Again, the continuum renormalization required for this procedure was at most a few percent in the spectral region of interest.
### Table 2.5: Lines used for the EMD reconstruction\(^a\). Method 2, with MEKAL/SPEX.

<table>
<thead>
<tr>
<th>Line</th>
<th>(\log T^o_{m}) (K)</th>
<th>(\lambda) (Å)</th>
<th>47 Cas B</th>
<th>EK Dra</th>
<th>(\pi^1) UMa</th>
<th>(\chi^1) Ori</th>
<th>(\kappa^1) Cet</th>
<th>(\beta) Com</th>
</tr>
</thead>
<tbody>
<tr>
<td>Fe XVII</td>
<td>6.7</td>
<td>15.01</td>
<td>483.2±17.2</td>
<td>376.9±13.9</td>
<td>61.5±2.4</td>
<td>52.0±1.6</td>
<td>43.5±1.4</td>
<td>8.5±0.5</td>
</tr>
<tr>
<td>Fe XVII</td>
<td>6.7</td>
<td>16.78</td>
<td>217.5±12.3</td>
<td>152.7±9.4</td>
<td>29.3±1.8</td>
<td>24.1±1.1</td>
<td>24.8±1.1</td>
<td>4.3±0.4</td>
</tr>
<tr>
<td>Fe XVIII</td>
<td>6.8</td>
<td>14.20</td>
<td>292.9±15.2</td>
<td>207.4±11.7</td>
<td>20.5±1.6</td>
<td>17.0±1.0</td>
<td>16.2±1.0</td>
<td>1.4±0.3</td>
</tr>
<tr>
<td>Fe XIX</td>
<td>6.9</td>
<td>13.52</td>
<td>222.9±26.6</td>
<td>126.6±19.6</td>
<td>14.5±2.9</td>
<td>8.3±1.6</td>
<td>6.8±1.5</td>
<td>0.8±0.6</td>
</tr>
<tr>
<td>Fe XX</td>
<td>7.0</td>
<td>12.83</td>
<td>289.7±21.9</td>
<td>162.5±59.0</td>
<td>7.4±1.8</td>
<td>4.3±1.1</td>
<td>5.8±1.0</td>
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</tr>
<tr>
<td>Fe XXI</td>
<td>7.0</td>
<td>12.29</td>
<td>210.2±24.7</td>
<td>149.1±16.3</td>
<td>8.3±2.1</td>
<td>7.3±1.3</td>
<td>8.7±1.4</td>
<td>–</td>
</tr>
<tr>
<td>Fe XXII</td>
<td>7.2</td>
<td>11.74</td>
<td>183.3±21.5</td>
<td>116.5±15.1</td>
<td>–</td>
<td>–</td>
<td>–</td>
<td>–</td>
</tr>
<tr>
<td>Fe XXIV</td>
<td>7.3</td>
<td>10.62</td>
<td>95.0±40.1</td>
<td>55.8±13.1</td>
<td>–</td>
<td>–</td>
<td>–</td>
<td>–</td>
</tr>
<tr>
<td>Fe XXV</td>
<td>7.8</td>
<td>1.85</td>
<td>68.0±17.8</td>
<td>22.6±81.0</td>
<td>–</td>
<td>–</td>
<td>–</td>
<td>–</td>
</tr>
<tr>
<td>O VIII</td>
<td>6.5</td>
<td>18.97</td>
<td>639.8±17.6</td>
<td>298.5±11.5</td>
<td>32.2±1.7</td>
<td>28.2±1.1</td>
<td>29.5±1.1</td>
<td>5.6±0.4</td>
</tr>
<tr>
<td>O VII</td>
<td>6.3</td>
<td>21.60</td>
<td>90.2±11.8</td>
<td>50.1±8.2</td>
<td>11.6±1.6</td>
<td>8.6±1.0</td>
<td>10.1±0.96</td>
<td>2.4±0.5</td>
</tr>
<tr>
<td>O VII</td>
<td>6.3</td>
<td>22.10</td>
<td>48.6±10.2</td>
<td>39.5±7.7</td>
<td>8.4±1.5</td>
<td>6.5±0.8</td>
<td>6.1±0.9</td>
<td>3.3±0.5</td>
</tr>
<tr>
<td>C VI</td>
<td>6.1</td>
<td>33.73</td>
<td>68.2±7.9</td>
<td>38.1±5.2</td>
<td>2.6±2.3</td>
<td>3.6±0.5</td>
<td>4.1±0.6</td>
<td>1.3±0.3</td>
</tr>
<tr>
<td>N VII</td>
<td>6.3</td>
<td>24.77</td>
<td>62.8±8.9</td>
<td>32.3±5.1</td>
<td>3.8±1.5</td>
<td>1.8±0.4</td>
<td>2.5±0.4</td>
<td>0.4±0.3</td>
</tr>
<tr>
<td>Ne X</td>
<td>6.8</td>
<td>12.13</td>
<td>475.0±33.3</td>
<td>185.7±21.2</td>
<td>–</td>
<td>–</td>
<td>–</td>
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</tr>
<tr>
<td>Ne IX</td>
<td>6.6</td>
<td>13.45</td>
<td>277.8±27.6</td>
<td>113.0±17.9</td>
<td>8.4±2.3</td>
<td>8.9±1.5</td>
<td>10.0±1.5</td>
<td>0.9±0.5</td>
</tr>
<tr>
<td>Ne IX</td>
<td>6.6</td>
<td>13.70</td>
<td>–</td>
<td>–</td>
<td>5.9±2.0</td>
<td>5.2±1.2</td>
<td>6.4±1.2</td>
<td>–</td>
</tr>
<tr>
<td>Mg XII</td>
<td>7.0</td>
<td>8.42</td>
<td>238.7±13.2</td>
<td>93.9±8.1</td>
<td>3.9±0.7</td>
<td>2.5±0.4</td>
<td>6.1±0.5</td>
<td>0.3±0.2</td>
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<tr>
<td>Mg XI</td>
<td>6.8</td>
<td>9.17</td>
<td>320.6±10.8</td>
<td>157.2±7.1</td>
<td>15.0±0.7</td>
<td>10.4±0.4</td>
<td>14.7±0.5</td>
<td>1.5±0.1</td>
</tr>
<tr>
<td>Si XIV</td>
<td>7.2</td>
<td>6.18</td>
<td>114.2±12.0</td>
<td>57.3±7.9</td>
<td>–</td>
<td>–</td>
<td>–</td>
<td>–</td>
</tr>
<tr>
<td>Si XIII</td>
<td>7.0</td>
<td>6.65</td>
<td>194.0±11.0</td>
<td>105.6±7.5</td>
<td>5.6±0.6</td>
<td>4.9±0.4</td>
<td>5.7±0.4</td>
<td>0.7±0.2</td>
</tr>
<tr>
<td>S XVI</td>
<td>7.2</td>
<td>4.72</td>
<td>55.6±12.0</td>
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<td>–</td>
<td>–</td>
<td>–</td>
<td>–</td>
</tr>
<tr>
<td>S XV</td>
<td>7.2</td>
<td>5.04</td>
<td>93.3±11.3</td>
<td>35.7±6.3</td>
<td>–</td>
<td>–</td>
<td>–</td>
<td>–</td>
</tr>
</tbody>
</table>

\(^a\) For each line, the measured luminosity in \(10^{26}\) erg s\(^{-1}\) using the MEKAL database is given. Note that the entries for the Fe lines contain blends of Fe around the given lines.

\(^b\) Maximum line formation temperature.
Table 2.6: Lines used for the EMD reconstruction\textsuperscript{a}. Method 2, with APEC/XSPEC

<table>
<thead>
<tr>
<th>Line</th>
<th>( \log T^\text{eff}_\lambda ) (K)</th>
<th>( \lambda ) (Å)</th>
<th>47 Cas B</th>
<th>EK Dra</th>
<th>( \chi^1 ) UMa</th>
<th>( \chi^1 ) Ori</th>
<th>( \kappa^1 ) Cet</th>
<th>( \beta ) Com</th>
</tr>
</thead>
<tbody>
<tr>
<td>Fe \textit{vii}</td>
<td>6.7</td>
<td>15.01</td>
<td>477.5±19.6</td>
<td>378.7±16.3</td>
<td>61.5±2.9</td>
<td>51.9±1.7</td>
<td>44.2±1.8</td>
<td>8.1±0.7</td>
</tr>
<tr>
<td>Fe \textit{vii}</td>
<td>6.7</td>
<td>16.78</td>
<td>222.3±18.6</td>
<td>150.5±15.2</td>
<td>30.9±2.7</td>
<td>25.1±1.8</td>
<td>25.5±1.6</td>
<td>4.5±0.7</td>
</tr>
<tr>
<td>Fe \textit{viii}</td>
<td>6.9</td>
<td>14.20</td>
<td>297.4±18.9</td>
<td>207.1±15.9</td>
<td>20.4±2.7</td>
<td>17.2±1.3</td>
<td>16.9±1.3</td>
<td>1.4±0.6</td>
</tr>
<tr>
<td>Fe \textit{ix}</td>
<td>6.9</td>
<td>13.52</td>
<td>235.2±39.8</td>
<td>114.7±34.0</td>
<td>14.4±2.3</td>
<td>5.8±2.8</td>
<td>6.6±2.2</td>
<td>0.7±0.6</td>
</tr>
<tr>
<td>Fe \textit{x}</td>
<td>7.0</td>
<td>12.83</td>
<td>246.8±30.9</td>
<td>144.7±30.1</td>
<td>8.6±1.8</td>
<td>4.4±1.9</td>
<td>7.1±1.5</td>
<td>–</td>
</tr>
<tr>
<td>Fe \textit{xx}</td>
<td>7.0</td>
<td>12.29</td>
<td>190.0±40.3</td>
<td>132.2±29.6</td>
<td>12.3±3.7</td>
<td>6.8±2.3</td>
<td>9.7±1.5</td>
<td>–</td>
</tr>
<tr>
<td>Fe \textit{xxii}</td>
<td>7.2</td>
<td>11.74</td>
<td>155.4±20.4</td>
<td>109.6±14.3</td>
<td>–</td>
<td>–</td>
<td>–</td>
<td>–</td>
</tr>
<tr>
<td>Fe \textit{xxiv}</td>
<td>7.3</td>
<td>10.62</td>
<td>71.3±59.4</td>
<td>53.5±33.5</td>
<td>–</td>
<td>–</td>
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<td>–</td>
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<tr>
<td>Fe \textit{xxv}</td>
<td>7.8</td>
<td>1.85</td>
<td>72.7±17.9</td>
<td>40.6±14.1</td>
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<tr>
<td>O \textit{viii}</td>
<td>6.5</td>
<td>18.97</td>
<td>651.1±17.7</td>
<td>291.4±11.4</td>
<td>32.9±1.7</td>
<td>28.4±1.1</td>
<td>30.1±8.7</td>
<td>5.6±0.4</td>
</tr>
<tr>
<td>O \textit{vii}</td>
<td>6.3</td>
<td>21.60</td>
<td>99.7±20.8</td>
<td>45.5±13.3</td>
<td>11.8±1.6</td>
<td>9.1±1.4</td>
<td>10.9±1.4</td>
<td>2.8±0.9</td>
</tr>
<tr>
<td>O \textit{vii}</td>
<td>6.3</td>
<td>22.10</td>
<td>43.8±10.6</td>
<td>39.5±11.8</td>
<td>8.4±1.5</td>
<td>6.4±1.2</td>
<td>6.5±1.2</td>
<td>3.1±0.9</td>
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<tr>
<td>C \textit{vi}</td>
<td>6.1</td>
<td>33.73</td>
<td>56.9±7.6</td>
<td>30.1±5.0</td>
<td>2.5±1.0</td>
<td>2.8±0.5</td>
<td>3.5±0.5</td>
<td>1.3±0.3</td>
</tr>
<tr>
<td>N \textit{vii}</td>
<td>6.3</td>
<td>24.77</td>
<td>69.2±9.2</td>
<td>32.4±5.2</td>
<td>2.1±0.9</td>
<td>1.5±0.4</td>
<td>2.7±0.5</td>
<td>0.2±0.2</td>
</tr>
<tr>
<td>Ne \textit{x}</td>
<td>6.8</td>
<td>12.13</td>
<td>500.2±35.3</td>
<td>167.0±21.7</td>
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<tr>
<td>Ne \textit{ix}</td>
<td>6.6</td>
<td>13.45</td>
<td>285.0±42.7</td>
<td>87.5±36.2</td>
<td>7.9±4.0</td>
<td>10.2±2.7</td>
<td>10.6±2.9</td>
<td>–</td>
</tr>
<tr>
<td>Ne \textit{ix}</td>
<td>6.6</td>
<td>13.70</td>
<td>–</td>
<td>–</td>
<td>5.1±2.0</td>
<td>6.0±1.3</td>
<td>7.1±1.2</td>
<td>–</td>
</tr>
<tr>
<td>Mg \textit{xii}</td>
<td>7.0</td>
<td>8.42</td>
<td>245.9±13.6</td>
<td>108.9±16.1</td>
<td>3.4±0.9</td>
<td>2.1±0.7</td>
<td>5.5±0.8</td>
<td>0.3±0.2</td>
</tr>
<tr>
<td>Mg \textit{xii}</td>
<td>7.0</td>
<td>9.17</td>
<td>319.5±11.3</td>
<td>167.5±13.5</td>
<td>12.9±1.4</td>
<td>9.6±0.7</td>
<td>13.2±0.8</td>
<td>1.5±0.2</td>
</tr>
<tr>
<td>Si \textit{xiv}</td>
<td>7.2</td>
<td>6.18</td>
<td>130.6±24.7</td>
<td>63.0±11.9</td>
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<tr>
<td>Si \textit{xiii}</td>
<td>7.0</td>
<td>6.65</td>
<td>207.3±75.0</td>
<td>104.5±14.0</td>
<td>6.5±1.2</td>
<td>4.5±0.6</td>
<td>5.5±0.9</td>
<td>0.7±0.1</td>
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<tr>
<td>S \textit{xvi}</td>
<td>7.4</td>
<td>4.72</td>
<td>52.9±8.1</td>
<td>–</td>
<td>–</td>
<td>–</td>
<td>–</td>
<td>–</td>
</tr>
<tr>
<td>S \textit{xv}</td>
<td>7.2</td>
<td>5.04</td>
<td>105.1±18.9</td>
<td>34.1±6.4</td>
<td>–</td>
<td>–</td>
<td>–</td>
<td>–</td>
</tr>
</tbody>
</table>

\textsuperscript{a} For each line, the measured luminosity is given in \( 10^{26} \text{ erg s}^{-1} \) using the APEC database. Note that the entries for the Fe lines contain blends of Fe around the given lines. \textsuperscript{b} Maximum line formation temperature. \textsuperscript{c} For \( \beta \) Com none of the Ne lines could be reliably measured by fitting \( \delta \) lines using the baseline model from method 1 and the APEC database within XSPEC.
To actually extract the line fluxes, we added $\delta$ function lines to the model, located at the theoretical wavelengths of the principal lines in the blends under consideration, and convolved them with the instrument response. The line fluxes were then obtained by fitting the amplitudes of the $\delta$ functions to the observed lines and the adjacent narrow wavelength ranges. If the wings of two Fe lines overlapped significantly, we fitted two $\delta$ functions simultaneously. Note that the fluxes thus derived may be different for APEC and MEKAL because of slightly different atomic physics, fit parameters, and blend contributions.

The obtained Fe, O\textsc{viii} and O\textsc{vii} line fluxes were converted to luminosities as presented in Table 2.5 and 2.6 for MEKAL and APEC emissivities, respectively, based on Hipparcos distances as quoted in Table 2.1. They were then used to reconstruct the EMD, as follows:

A first, smooth estimate of the EMD was derived from the emissivities at the maximum line formation temperature $T_m$ of each Fe line. The emissivities are based on solar abundances as given by Anders & Grevesse (1989). Our EMD was defined on a grid with a grid point separation of $\Delta \log T = 0.1$ dex in the range between $\log T = 5.5 - 8$ for SPEX ($T$ in K) (APEC: $\log T = 5.7 - 8$) for the more active stars 47 Cas B and EK Dra, and in the range between $\log T = 5.5 - 7.1$ (APEC: $\log T = 5.7 - 7.1$) for the less active stars $\chi^1$ Ori, $\kappa^1$ Cet and $\beta$ Com. For $\pi^1$ UMa, the EMD was defined in the range between $\log T = 5.5 - 7.3$ (APEC: $\log T = 5.7 - 7.3$). (The different low-temperature limits reflect the different availability of emissivities in the different codes, but we note that the range of $\log T = 5.5 - 5.7$ is irrelevant for any line we use for the EMD reconstruction, and it is also not important for the emissivities of any lines of other elements analyzed here.) As a starting condition, the EMD was extrapolated to temperatures cooler than $T_m$ of Fe\textsc{xvii} ($\log T = 6.7$) by using a slope of 2, and to temperatures higher than $T_m$ of the hottest Fe ion in use (Fe\textsc{xxv}, $\log T = 7.9$ for 47 Cas and EK Dra, Fe\textsc{xix}, $\log T = 6.9$ for $\beta$ Com, and Fe\textsc{xxi}, $\log T = 7.0$ for the other stars) by using a slope of $-2$. These extrapolations were performed to the limits of the respective temperature ranges as defined above. These starting conditions were suggested by the slopes of the initial EMDs defined by the EM values at the different $T_m$. Also, we subsequently found that the low-$T$ slope converges to values around 2 even if the starting slope was largely different (see Sect. 2.7.3). For each line, the flux $F_C$ predicted from the initial EMD was calculated according to the following equation:

$$F_C(x) = \frac{1}{4\pi d^2} \sum_i EM(i) \cdot \varphi(x,i)$$ (2.3)

where the sum is over the temperature bins $i$; $EM(i)$ is the emission measure in the $i$th temperature bin, and $\varphi(x,i)$ is the emissivity of the given line $x$ at this grid point. The calculated fluxes were then compared with the measured
2.5. Data Analysis

uxes, \( F(x) \). The EM in each bin was iteratively corrected using the algorithm described by Withbroe (1975):

\[
EM^{n+1}(i) = EM^n(i) \cdot \frac{\sum_x \frac{F(x)}{F_C(x)} \cdot \varphi(x, i) + \frac{RC}{R} \cdot \varphi(O \text{ VII}, i)}{\sum_x \varphi(x, i) + \varphi(O \text{ VII}, i)} \tag{2.4}
\]

where \( i \) is the index of the temperature bins and \( n \) is the iteration number.

The last term in the numerator and denominator was added to also obtain convergence of the line-ux ratio between \( O \text{ VIII} \) and \( O \text{ VII} \), where we have used the measured \( (R) \) and the predicted \( (R_C) \) ux ratio. The EMD was then iterated until we reached a pre-set convergence criterion.

In the presence of considerable systematic uncertainties in the modeling, a reasonable convergence criterion should be set, although its statistical meaning may be marginal. A straightforward goal is to achieve convergence in such a way that on average the squared deviation between model and fit for a line ux is equal to the variance of the same line ux. This leads to a reduced-\( \chi^2 \)-like expression,

\[
\psi^2 = \frac{1}{N_x + 1} \left( \sum_x \left[ \frac{(F[x] - F_C[x])^2}{(\sigma[x])^2} + \frac{(R - R_C)^2}{(\sigma_R)^2} \right] \right) \tag{2.5}
\]

where the last term relates to the \( O \text{ VIII}/O \text{ VII} \) flux ratio. Here, \( N_x \) is the number of lines used, and \( \sigma(x) \) is the error assigned to the measured line flux, which comprises the error from the finite photon statistics and an assumed systematic uncertainty from the atomic physics (see Sect. 2.5.3). Similarly, \( \sigma_R \) is the derived error in the \( O \text{ VIII}/O \text{ VII} \) flux ratio. We iterated until this expression reached a value of unity, or if this did not occur, until it no longer significantly decreased. We would like to mention that the fit parameters, i.e., the EMs per bin, are not independent of each other owing to the broad emissivity curves of each line; given the large systematic uncertainties in the line uxes, we feel that a more detailed convergence criterion is not warranted. Our main goal is to stop the iteration at a reasonable level to avoid over-interpretation of superficial features in the EMD that may arise from iterating too deeply - see our discussion in Sect. 2.6.1.

At this point, then, we have found an EMD from uxes of the single element Fe, under the assumption of solar metallicity as used for the tabulated emissivities, and from a ratio of O line-uxes. Therefore, there still remains a normalization factor for the EMD and the absolute level of the Fe abundance to be determined (see Sect. 2.5.3). We will use the observed continuum level to fix the EMD normalization and, at the same time, the absolute Fe abundance, as explained in the following.
Abundances

To determine the abundances of the elements, we extracted all lines of interest (see Tables 2.5 and 2.6) in a similar manner as we extracted the Fe and O lines before. Here, however we used the EMD constructed from our Fe and O line fluxes in order to describe the continuum, instead of the 10-\( T \) model. This EMD is known only up to a normalization constant depending on the absolute Fe abundance, which we derived as follows: we constructed a set of spectra from the calculated EMD with different Fe abundances such that the product of the Fe abundance and the EM(\( T \)) is constant for any \( T \), still using the approximate abundances of the other elements from method 1 to estimate the contributions from blends and to obtain a more accurate description of the continuum. From the spectrum that best fitted the nearly line-free regions long- and shortward of the O viii Ly\( \alpha \) line at 18.97 Å, we obtained the absolute Fe abundance. We then extracted the line fluxes of all interesting elements using \( \delta \)-line models, as described before. Note that the line fluxes of Mg, Si, S, and, if available, Fe xxv, were extracted from the MOS spectra while all other lines were extracted from the RGS spectra. As done for the RGS spectra before, we adjusted the model continuum to the observation in the line-free regions of the MOS instrument at high energies, before the line-flux measurement.

The predicted line fluxes of these elements (see Table 2.5 and 2.6) were then calculated from the EMD, using their catalogued emissivities, which were again based on solar photospheric abundances. The ratios between the predicted and the measured fluxes provided the abundances relative to Fe, \( A/A(\text{Fe}) \), with respect to the corresponding solar ratios. In some cases, we measured the fluxes of more than one line for a single element. In these cases, we calculated the abundance for each line and computed averages, using \( 1/\sigma_A^2 \) as weights, where \( \sigma_A \) is the error in the abundance (see Sect. 2.5.3). We implicitly assumed here that the abundance of an element is the same at all temperatures, an assumption that is not necessarily supported from solar observations (Jordan et al. 1998). The data quality at hand does not allow for further discrimination, however.

Finally, we once more iterated the adjustment of the continuum level to obtain the absolute Fe abundance, as described above, now using the abundances determined from our procedure. The updated values closely agreed with the previously obtained Fe abundances. The error of the absolute Fe abundance was derived by varying it around its best fit value and requiring that the continuum fit be acceptable within one sigma. The final abundance values we report in this chapter refer to the solar photospheric abundances given by Anders & Grevesse (1989), except for Fe for which we adopt the value given by Grevesse & Sauval (1999).
2.5. Data Analysis

Errors

Errors arise from different sources. First, they are due to uncertainties in the atomic databases; these are not easily quantifiable, but are likely to be in the range of several percent to perhaps 20%, depending on the line under consideration. For method 2, we chose mostly bright, well-studied lines, and for the sake of definition we have assumed systematic uncertainties of 10% for each line. Statistical errors also arise from the fit of the \( \delta \)-line model used to extract the line flux; these essentially originate from photon statistics. For each line flux, these two errors were summed in quadrature, and we call them line-errors \( \sigma \).

The errors in the EMD were estimated by statistically varying the fluxes of the line blends according to their \( \sigma \), and repeating the EMD reconstruction for 19 different, perturbed line-flux lists. We thus derived formal upper and lower 1-sigma ranges of the EMD solutions by using standard formulae from Gaussian statistics, although we mention that the various solutions are not necessarily normally-distributed in \( \log \) EM at any given temperature. The standard deviations thus derived, however, provide a well-defined characteristic width of the distribution, and we verified that the ranges containing 68\% of the solutions and the 1-sigma ranges are very similar. Because the EM values scatter considerably in any temperature bin, we performed the statistics using logarithms of the EMs in each temperature bin in order to avoid the average being biased by one or a few large values. We extended the error analysis to include up to 100 perturbed line lists but the error ranges did not significantly change. We caution that the EMD slopes on both sides of the peak temperature could be slightly dependent on the initial EMD guess, where we assume a slope of \( \pm 2 \) (see Sect. 2.5.3). We study the EMD results starting from different initial conditions for the slopes in Sect. 2.7.3. The final EMD slopes converged to similar values. These effects are not taken into account in the EMD errors reported here. We note that these errors are only given as an indicator for the uncertainty in the EMD, but they are not explicitly used further in our error analysis.

The error of the abundance \( A \) is proportional to the line-error, \( \sigma_A = A \cdot \sigma(x)/F(x) \). If the abundance is a weighted average, then the final error is the larger of i) \( \sum 1/\sigma_A^2 \)^{-1/2} and ii) the error of the weighted means of all abundance values. Moreover, there is an error in the abundances arising from the variation in the 20 different EMD reconstructions: for each of these reconstructions, we found slightly different abundances with new errors. We defined the error from this variation as the larger of i) the average error found in each reconstruction, and ii) the standard deviation of the twenty abundance values per element. We note, however, that we adopted the abundance derived from the best-fit solution, with no perturbation applied.

Finally, the error in the adjustment of the continuum (required to deter-
mine the line fluxes, Sect. 2.5.3) also affects the abundance errors. As our final error for a given abundance, we summed in quadrature the error related to the line-error (or the average if multiple lines were used, as defined above), the error arising from the variation of the EMD, and the error from the continuum adjustment. We emphasize, however, that this procedure can provide no more than a simulated estimate of realistic errors. The unknown systematic deviations in the atomic physics parameters prevent us from obtaining better estimates.

2.6 Results

2.6.1 Emission Measure Distributions

The EMDs derived from the two different methods are shown in Figure 2.5. In the left column, EMDs from method 1 using SPEX in combination with a fit based on Chebychev polynomials of degree of 6 and, where possible, 8 are plotted. The middle and the right columns show EMDs reconstructed with our method 2, based on MEKAL and APEC emissivities, respectively. In the middle and right columns, the black histograms illustrate the best-fit EMDs while the red histograms mark the 1σ range at each temperature, derived from the perturbed flux lists. As the best-fit EMDs are derived from unperturbed fluxes, and they are not equal to the mean EMDs derived from the perturbed flux lists, these ranges of variation do not need to be symmetrically arranged around the best-fit solutions. In some cases (EK Dra and 47 Cas), the lower error ranges drop rapidly to very low values at certain temperatures. Although this is a consequence of the increasing uncertainty in the EMD at the lowest and the highest temperatures, we note that the error ranges are given on a logarithmic scale; once the ranges becomes large, the precise level of the lower bound is of little importance.

For method 2 the quality of the EMD can be measured by comparing the predicted and the observed line fluxes. The final agreement between predicted and observed line fluxes is illustrated in Fig. 2.6, where we show the fractional deviation of the predicted line fluxes from the observed values, \( (F_{\text{C}} - F)/F \) for Fe and O VIII/O VII flux ratio. Most line fluxes agree within 10–20%, with the larger deviations mainly relating to the weakest lines, i.e., the lines formed at high temperatures in the least active stars (e.g., Fe XX for \( \pi^1 \) UMa).

The EMDs derived from the different methods show rather similar characteristics. We see that the temperature where the EMD peaks decreases toward older, less active stars, namely from about 10 MK for 47 Cas B and EK Dra to 5 MK for \( \pi^1 \) UMa, \( \chi^1 \) Ori and \( \kappa^1 \) Cet, and to \( \lesssim 4 \) MK for our oldest target, \( \beta \) Com. Characteristic values for the Sun are 1–3 MK, depending on the phase of its activity cycle (Peres et al. 2000). The average temperatures derived from
2.6. Results

Figure 2.5: Reconstructed emission measure distributions. **Left:** From method 1 using SPEX, based on Chebychev polynomials of order 6 (solid line) and order 8 (dashed). **Middle:** From method 2, based on MEKAL emissivities. **Right:** From method 2, based on APEC emissivities. In the middle and right plots, the red histograms illustrate the $\pm 1\sigma$ range of solutions from the average of 20 EMDs reconstructed from the original and from the perturbed line lists. The black histograms illustrate the best-fit EMDs, derived from the unperturbed line-flux list.
method 2 (using SPEX), log $T$, are listed in Table 2.7. To obtain these values, we calculated the mean of log $T$, using the EMs in each bin as weights. Also given are the lower and upper threshold temperatures that comprise 90% of the total EM (on each side of the EMD peak).

With method 2, we generally obtain a smoother and flatter EMD than with method 1 and the Chebychev polynomial approximation. At temperatures above log $T \approx 7.5$, the EMD is not well constrained. This is obvious for the cooler coronae which do not provide any useful spectral lines at those temperatures, and we therefore did not extend our DEM analysis to this range. Considerable scatter is still also found for the EMDs of 47 Cas B and EK Dra, despite the availability of Fe XXIII-Fe XXV lines. The reason resides in the fact that Fe XXIII and Fe XXIV show very faint lines, and Fe XXV is the only blend complex that covers the temperatures above log $T \approx 7.5$. In the more active stars, the EMDs from method 1 seem to be composed of two peaks at about 6 and 20 MK, well separated by a local minimum. Also, in most EMDs derived from method 1, we find a deep decrease in the EM below about log $T \approx 6.2$, combined with a local EM peak around log $T = 6.0$. Both features become stronger if higher polynomial degrees are used.

Several effects may contribute to this: First, a large range of solutions may in fact be compatible with the spectra, given that the spectral inversion
Table 2.7: Mean coronal temperatures\(^a\) and ranges

<table>
<thead>
<tr>
<th>Star</th>
<th>log (L^L_X)</th>
<th>(\log \bar{T})</th>
<th>(\log T_{\text{min}})</th>
<th>(\log T_{\text{max}})</th>
<th>(\log \bar{T})</th>
<th>(\log T_{\text{min}})</th>
<th>(\log T_{\text{max}})</th>
</tr>
</thead>
<tbody>
<tr>
<td>47 Cas B</td>
<td>30.39</td>
<td>7.03</td>
<td>6.51</td>
<td>7.60</td>
<td>7.02</td>
<td>6.52</td>
<td>7.65</td>
</tr>
<tr>
<td>EK Dra</td>
<td>30.08</td>
<td>6.96</td>
<td>6.46</td>
<td>7.45</td>
<td>6.99</td>
<td>6.47</td>
<td>7.58</td>
</tr>
<tr>
<td>π(^1) UMa</td>
<td>29.06</td>
<td>6.65</td>
<td>6.16</td>
<td>6.98</td>
<td>6.69</td>
<td>6.18</td>
<td>7.17</td>
</tr>
<tr>
<td>(\chi^1) Ori</td>
<td>28.95</td>
<td>6.64</td>
<td>6.14</td>
<td>7.03</td>
<td>6.66</td>
<td>6.22</td>
<td>6.98</td>
</tr>
<tr>
<td>(\kappa^1) Cet</td>
<td>28.95</td>
<td>6.66</td>
<td>6.18</td>
<td>7.05</td>
<td>6.65</td>
<td>6.13</td>
<td>7.10</td>
</tr>
<tr>
<td>(\beta) Com</td>
<td>28.26</td>
<td>6.59</td>
<td>6.15</td>
<td>6.92</td>
<td>6.55</td>
<td>6.07</td>
<td>6.89</td>
</tr>
</tbody>
</table>

\(^a\) \(\log \bar{T}\) is the EM-weighted average of \(\log T\), \(T_{\text{min}}\) and \(T_{\text{max}}\) are the minimum and maximum temperatures that contain 90% of the EM above and below \(T\), respectively.

\(^b\) Luminosity in erg s\(^{-1}\) determined with \textit{XMM-Newton} in the 0.1–10 keV band.

is ill-conditioned, in particular in temperature regions where few constraints are available. Second, the fit of method 1 iterates to minimum \(\chi^2\), which considers only the Poissonian errors in the count spectrum and which may introduce EMD features of little relevance given the systematic uncertainties in the atomic physics, while method 2 has been terminated according to \(\psi^2\) (see Eq. 2.5), which approximately considers the atomic physics uncertainties as well. Third, method 1 uses many lines that may introduce uncertainty to the spectral fit, while the result of method 2 almost uniquely relies on the \(\text{O viii}/\text{O vii}\) flux ratio for the coolest portion of the EMD. And finally, method 1 imposes polynomial constraints on the solution, which favors the appearance of peaks and valleys in the EMD, while method 2 starts with a smooth EMD that is changed only in so far as the spectrum requires. For example, if a line requires excess EM due to an underestimation of its emissivity at a given temperature, then the reconstruction process may compensate by lowering the EM at adjacent temperatures as dictated by lines dominating there. To test this hypothesis, we iterated method 2 excessively, to \(\psi^2 = 0.5\). As the two examples in Fig. 2.7 show, very similar features also evolve in these examples. The low-temperature slopes appear to become partly steeper as well.

We also note that the amplitudes of the oscillations are compatible with the error ranges from perturbing the line-flux lists used in method 2 (cf. Fig. 2.5). It is conceivable that the oscillations found after a deep iteration of method 2 correspond to those seen in method 1 although this cannot be explicitly proven, given the largely different approaches. However, the magnitudes of the oscillations appear to be similar. It is also possible that the oscillations are present in the stellar EMD; we cannot reliably discriminate between this
Figure 2.7: Emission measure distributions (using SPEX with method 2).  
Left: The iteration was stopped when \( \psi^2 = 1 \) was reached. Right: The iteration was stopped when \( \psi^2 = 0.5 \) was reached. Note the additional oscillations in the EMD.

hypothesis and a numerical effect as long as we include statistical errors and assume the presence of systematic uncertainties of the magnitude adopted here (see Sect. 2.5.3).

In contrast, the abundance ratios turn out to be robust, with no significant change when deeper iterations are applied. In fact, the synthesized spectra for the two cases are very similar, i.e., the two EMDs represent the spectra almost equally well. Comparing the synthesized spectrum with the observations in the wavelength intervals illustrated in Fig. 2.3, we find, for 47 Cas B, a reduced \( \chi^2 \), \( \chi^2_{\text{red}} = 1.28 \) (for 1091 d.o.f) for the deeper integration and \( \chi^2_{\text{red}} = 1.37 \) (for 1091 d.o.f) for our standard convergence criterion. Although at first sight appearing significant, the important line features and the continuum in the RGS are fitted well in both cases.

We note that the EPIC MOS portion of the synthetic spectrum from method 2 is not very well fitted: the fit lies systematically below the data. This feature is probably to be ascribed to a cross-calibration inaccuracy in the effective areas of the RGS and the EPIC MOS instruments. As a matter of fact, a related effect is present in the spectrum from method 1 as well, but there, the continuum level of the synthetic spectrum is slightly but systematically too high in the RGS compared to the data, while the fit in the MOS spectrum is better. This is one important source for the somewhat higher \( \chi^2_{\text{red}} \) for our method 2. We note, however, that the analysis based on method 2 corrects for the continuum discrepancy before line-flux extraction by adjusting
Table 2.8: Reduced $\chi^2$ for the synthetic best-fit spectra$^a$

<table>
<thead>
<tr>
<th>Star</th>
<th>$\chi^2_{\text{red}}$ $^b$</th>
<th>$\chi^2_{\text{red}}$ $^c$</th>
<th>$\chi^2_{\text{red}}$ $^d$</th>
<th>d.o.f $^e$</th>
</tr>
</thead>
<tbody>
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<td>47 Cas B</td>
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<td>1.37</td>
<td>1.28</td>
<td>1091</td>
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<td></td>
<td>1.14</td>
<td>1.35</td>
<td>1006</td>
<td></td>
</tr>
<tr>
<td>EK Dra</td>
<td>1.25</td>
<td>1.34</td>
<td>1.25</td>
<td>769</td>
</tr>
<tr>
<td></td>
<td>1.21</td>
<td>1.29</td>
<td>671</td>
<td></td>
</tr>
<tr>
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<td>1.66</td>
<td>1.51</td>
<td>1.25</td>
<td>415</td>
</tr>
<tr>
<td></td>
<td>1.28</td>
<td>1.42</td>
<td>327</td>
<td></td>
</tr>
<tr>
<td>$\chi^1$ Ori</td>
<td>1.58</td>
<td>1.71</td>
<td></td>
<td>628</td>
</tr>
<tr>
<td></td>
<td>1.28</td>
<td>1.39</td>
<td>514</td>
<td></td>
</tr>
<tr>
<td>$\kappa^1$ Cet</td>
<td>1.40</td>
<td>1.58</td>
<td></td>
<td>635</td>
</tr>
<tr>
<td></td>
<td>1.35</td>
<td>1.43</td>
<td>538</td>
<td></td>
</tr>
<tr>
<td>$\beta$ Com</td>
<td>1.33</td>
<td>1.42</td>
<td></td>
<td>405</td>
</tr>
<tr>
<td></td>
<td>1.24</td>
<td>1.27</td>
<td>313</td>
<td></td>
</tr>
</tbody>
</table>

$^a$ The first line for a given star is based on MEKAL emissivities, the second line on APEC emissivities.

$^b$ Spectrum obtained from method 1 EMD, using the regions listed in Table 2.4.

$^c$ Spectrum obtained from method 2 EMD, iterated to $\psi^2=1.0$ (based on MEKAL).

$^d$ Spectrum obtained from method 2 EMD, iterated to $\psi^2=0.5$ (based on MEKAL).

$^e$ Degrees of freedom.

the continuum level individually both for the RGS and the EPIC spectra (see Sect. 2.5.3).

In Table 2.8, we compare the $\chi^2_{\text{red}}$ values with respect to the observations, for the synthetic spectra constructed from the EMDs that were obtained with method 1 and with method 2, respectively. Again, only the regions listed in Table 2.4 are used. The degrees of freedom are also listed in Table 2.8. Although formally the same wavelength intervals were used, the number of degrees of freedom are somewhat different in SPEX and XSPEC. This discrepancy comes from two different sources: first, partial bins at the beginning and the end of each interval are considered differently (XSPEC ignoring all partial bins). Although this discrepancy could be reduced by minor adjustments, we prefer to keep with the simple prescriptions of Table 2.4 for easy reproduction of our results. The partial bins occur in relatively low-flux, shallow regions of the spectrum where essentially continuum is fitted, hence a difference by a single, usually well fitted continuum bin is of little relevance. Second, the standard spectral software packages treat grouped bins that may contain bad data channels differently (SPEX breaks bins up into partial bins, while XSPEC does not). This is a feature of the standard software packages that we test here.
The $\chi^2_{\text{red}}$ from method 2 are slightly larger but the differences are not very substantial despite the systematic uncertainties in the emissivities adopted in method 2 but not in method 1. Beside the fact that in the reconstructed spectrum of method 1 the MOS spectrum is better fitted (given its higher S/N ratio), the slightly better $\chi^2_{\text{red}}$ for method 1 could also be a result of excessively deep iterations that aim at fitting poorly-described line fluxes at the cost of smoothness in the EMD. It is noteworthy that the $\chi^2_{\text{red}}$ values from the deep iterations in method 2 closely approach the $\chi^2_{\text{red}}$ values of method 1.

There is ample literature on EMDs of active stars available. Several other authors have also found EMDs with features resembling ours. In particular, double-peaked EMDs have previously been reported, e.g., by Mewe et al. (1996) for AB Dor, Kaastra et al. (1996a) for RS CVn binaries, Güdel et al. (1997a) for solar analogs, Sanz-Forcada et al. (2001) for the RS CVn binary $\lambda$ And, and Huenemoerder et al. (2003) for AR Lac, using entirely different EMD reconstruction methods. In the light of our discussion on errors and iteration depths in Sect. 2.5 and 2.6, we cannot be certain on the actual reality of bi-modal DEMs, and given that other authors use similar atomic physics databases, the same caveat may apply to other EMD reconstructions as well. Güdel et al. (1997a) argued, based on solar-flare data, that a bi-modal structure can arise from the rapid decay of the EM of a population of flares as they are cooling. The present quality of our EMD inversions does not allow us to make more definitive conclusions at this point.

2.6.2 Abundances

The abundances found with the different methods are listed in Table 2.9 and plotted as a function of the FIP in Figure 2.8. We plot the abundance ratios $A/\text{Fe}$ with respect to the solar photospheric ratios as a function of the FIP. The open circles represent the coronal abundances derived from method 1 whereas the filled circles show the abundances derived from method 2. We find a good agreement between the abundance sets, and an acceptable overall agreement between the results using the MEKAL database and the APEC database.

Nevertheless, some differences can be noted in Table 2.9. In the older stars $\chi^1$ Ori, $\kappa^1$ Cet, and $\beta$ Com, some systematic differences occur for the C/Fe abundance ratios, the abundances derived with APEC being smaller than those derived with SPEX although the error ranges are large. For the same stars, some differences are also present in the N/Fe abundance ratio. In both cases, the abundances are determined from only one faint line ($\text{C}\,\text{vi}$ and $\text{N}\,\text{vii}$, respectively), making the measurements of the line fluxes difficult (see Tables 2.5 and 2.6). Differences also occur in the four older stars for the Ne/Fe abundance ratio. This is partly due to the lack of strong and reliable (i.e., unblended) Ne lines. While in hot coronae, the strong NeX line serves as a reliable indicator for the Ne abundance, this line is much weaker in cooler
2.6. Results

<table>
<thead>
<tr>
<th>Abun.</th>
<th>Method</th>
<th>47 Cas B</th>
<th>EK Dra</th>
<th>π1 UMa</th>
<th>χ^2</th>
<th>Cet</th>
<th>β Com</th>
</tr>
</thead>
<tbody>
<tr>
<td>C/Fe</td>
<td>M1 MEK</td>
<td>0.95±0.12</td>
<td>0.54±0.14</td>
<td>0.20±0.07</td>
<td>0.29±0.09</td>
<td>0.34±0.08</td>
<td>0.29±0.09</td>
</tr>
<tr>
<td></td>
<td>M2 MEK</td>
<td>0.80±0.18</td>
<td>0.63±0.19</td>
<td>0.19±0.24</td>
<td>0.34±0.11</td>
<td>0.43±0.12</td>
<td>0.66±0.29</td>
</tr>
<tr>
<td></td>
<td>M1 APEC</td>
<td>0.76±0.13</td>
<td>0.60±0.20</td>
<td>0.23±0.13</td>
<td>0.19±0.11</td>
<td>0.22±0.08</td>
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</tr>
<tr>
<td></td>
<td>M2 APEC</td>
<td>0.65±0.17</td>
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<td>0.14±0.09</td>
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<td>0.24±0.08</td>
<td>0.39±0.22</td>
</tr>
<tr>
<td>N/Fe</td>
<td>M1 MEK</td>
<td>1.18±0.21</td>
<td>0.84±0.18</td>
<td>0.36±0.15</td>
<td>0.25±0.07</td>
<td>0.43±0.16</td>
<td>0.12±0.15</td>
</tr>
<tr>
<td></td>
<td>M2 MEK</td>
<td>1.07±0.28</td>
<td>0.78±0.25</td>
<td>0.44±0.28</td>
<td>0.26±0.09</td>
<td>0.42±0.12</td>
<td>0.30±0.29</td>
</tr>
<tr>
<td></td>
<td>M1 APEC</td>
<td>1.20±0.21</td>
<td>0.71±0.21</td>
<td>0.32±0.14</td>
<td>0.18±0.04</td>
<td>0.24±0.08</td>
<td>0.03±0.06</td>
</tr>
<tr>
<td></td>
<td>M2 APEC</td>
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<td>0.73±0.21</td>
<td>0.20±0.14</td>
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<tr>
<td>O/Fe</td>
<td>M1 MEK</td>
<td>0.79±0.10</td>
<td>0.60±0.27</td>
<td>0.27±0.06</td>
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<td>M2 MEK</td>
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<td>M1 APEC</td>
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<td>Ne/Fe</td>
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<td>M2 MEK</td>
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<td>0.96±0.20</td>
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<td>0.32±0.09</td>
<td>0.44±0.10</td>
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<td>Mg/Fe</td>
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<tr>
<td>Si/Fe</td>
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<td>–</td>
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<td>–</td>
<td>–</td>
<td>–</td>
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<tr>
<td>Ar/Fe</td>
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<td>1.58±0.43</td>
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<td>–</td>
<td>–</td>
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<td>M1 APEC</td>
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<td>–</td>
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<tr>
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<td>M2 MEK</td>
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<td>0.87±0.14</td>
<td>1.18±0.22</td>
<td>1.27±0.33</td>
</tr>
<tr>
<td></td>
<td>M1 APEC</td>
<td>0.69±0.06</td>
<td>0.74±0.10</td>
<td>1.07±0.14</td>
<td>0.98±0.14</td>
<td>0.93±0.17</td>
<td>1.76±0.43</td>
</tr>
<tr>
<td></td>
<td>M2 APEC</td>
<td>0.55±0.05</td>
<td>0.96±0.11</td>
<td>1.26±0.31</td>
<td>0.83±0.13</td>
<td>1.83±0.48</td>
<td>–</td>
</tr>
</tbody>
</table>

*a All abundance ratios and Fe abundances are with respect to the solar photospheric abundances given by Anders & Grevesse (1989) except for Fe, for which the photospheric value given by Grevesse & Sauval (1999) has been adopted.

coronae and strongly blended with Fe lines. The Ne IX lines at 13.55–13.7 Å are always blended with Fe lines, but also become quite faint in the less active stars. Further, the inferred Ne flux in the line feature depends on the absolute Fe abundance. The latter is poorly determined in particular in the less active
Figure 2.8: Abundances relative to Fe as a function of FIP, normalized to solar photospheric ratios (Anders & Grevesse 1989; Grevesse & Sauval 1999). Open circles: method 1; filled circles: method 2.

stars where almost no continuum is present. We emphasize that we cannot attribute the discrepancy to any of the methods. The available data quality simply makes the determination of the Ne/Fe abundance ratio in cooler coronae ambiguous.

Fe blending leads to some differences in the Mg/Fe abundances as well (the largest deviations are for $\pi^1$ UMa using APEC method 2, and $\kappa^1$ Cet using SPEX method 2). Finally the weakness of the S lines in the more active stars leads to some differences in the S/Fe abundances.
2.7. Discussion

The absolute Fe abundance was systematically higher when the APEC database was used. Note that in the spectrum of $\beta$ Com, the continuum is almost nonexistent and the derivation of the Fe abundance is difficult. For the latter target, the $\delta$-fit did not converge for the Ne line with APEC. For this reason, these two points are missing in Table 2.9 and Figure 2.8.

However, we recognize many of these features to be due to limitations of the data and the reconstruction methods. On the other hand, these systematic differences are small compared to the general trends. As shown in Figure 2.8, the abundances resulting from the two methods and the two databases agree mostly quite well within the errors, and the general trends are the same, regardless of the method used.

2.6.3 Light Curves

The light curves of the six stars are shown in Figure 2.9. For each star, four light curves are plotted. They describe, from top to bottom, the total count rates in the 0.2 – 10 keV range (black), in the soft band (0.2 – 1 keV, green), in the hard band (> 1 keV), and the ratio between the hard and the soft count rates (blue). The upper energy thresholds used for the hard band vary from star to star and are listed in Table 2.4. Only data from detectors that operated in imaging mode were considered (for 47 Cas B, EK Dra and $\beta$ Com, the data from MOS1, MOS2 and PN were used; for $\pi^1$ UMa, data from MOS1 and MOS2, and for $\chi^1$ Ori and $\kappa^1$ Cet, only data from one MOS camera, MOS2 and MOS1 respectively, were used).

The light curves display considerable variability. We observe the presence of large flares on all stars in our sample except $\beta$ Com. However, even after excluding these flares, the light curves still show considerable variability that cannot be described by steady, quiescent coronal emission. We also note that the hard emission becomes weaker toward older stars, in agreement with the decline of the average coronal temperature described earlier.

2.7 Discussion

2.7.1 Correlation between the Parameters

The results of our analysis of six solar analogs clearly show a number of trends that we wish to quantify below on our results from method 2. Before doing so, we note that the X-ray results from the observation of $\kappa^1$ Cet are rather similar to results from $\pi^1$ UMa and $\chi^1$ Ori, despite the former's significantly longer rotation period and, hence, age. The reason for this discrepancy is not entirely clear, but we note that $\kappa^1$ Cet has a slightly later spectral type and therefore a somewhat lower mass than the other targets. It is known that later-type stars
Figure 2.9: Light curves of our targets. The four light curves in each panel show, from top to bottom, the total count rate in the $0.2 - 10$ keV range (black), in the soft band ($0.2 - 1$ keV, green), in the hard band ($> 1$ keV, red, where the upper energy limit is reported in the caption of Table 2.4), and the ratio of hard/soft (blue). For illustration purposes, the hardness ratio has been multiplied by 5, 3, 0.5, 0.5, 0.5, and 0.5 for 47 Cas B, EK Dra, π¹ UMa, χ¹ Ori, κ¹ Cet, and β Com, respectively. The bin size is, for the stars as listed above, 300, 300, 300, 450, 450, and 600 s, respectively. Only data from detectors that were operated in imaging mode were used, i.e., data from the PN camera were not used for π¹ UMa, χ¹ Ori, and κ¹ Cet. For the latter two stars, only one MOS camera was available in imaging mode (see Table 2.3). The time ranges of the largest flares excluded from the spectral analysis are also shown by dash-dotted vertical lines.

evolve more slowly (Soderblom et al. 1993) and that they remain in a state of maximum X-ray luminosity (the saturation limit) for longer rotation periods (Pizzolato et al. 2003) than stars of earlier spectral type. Both effects make
2.7. Discussion

Figure 2.10: Mean coronal temperature as a function of the X-ray luminosity. The dashed and solid lines are the regression fits to the results based on APEC and MEKAL, respectively.

κ1 Cet look somewhat younger than inferred from a rotation-activity relation that is appropriate for early G stars.

In Figure 2.10, we plot the mean coronal temperature, $T$, (see Table 2.7) as a function of the total luminosity $L_X$. We fitted the data with a power-law. Because both variables to be correlated are likely to be affected by systematic scatter around any power-law, we use the ordinary least squares bisector method as described by Isobe et al. (1990). Clearly, the two parameters are correlated. We find, for our results from MEKAL and APEC, respectively,

$$L_X \approx 1.17 \times 10^{26}T^{4.26\pm0.41} \text{ erg s}^{-1} \text{ (MEKAL)},$$  \hspace{1cm} (2.6)

$$L_X \approx 1.61 \times 10^{26}T^{4.05\pm0.25} \text{ erg s}^{-1} \text{ (APEC)},$$  \hspace{1cm} (2.7)

where $T$ is in MK. These relations are consistent with the results of Güdel et al. (1997a), except that the latter authors used the higher temperature for a model with two thermal components fitted to ROSAT spectra. Two points for the Sun at minimum and maximum activity level are also plotted for comparison (after Peres et al. 2000).

In Figure 2.11 the temperature as a function of the period is shown. Again, we fitted the data with a power-law. In this case, the relations calculated from
Figure 2.11: Mean coronal temperature as a function of the stellar rotation period. Open circles refer to the APEC values, filled circles to the MEKAL values, both based on method 2. The dashed and solid lines are the regression fits to the APEC and MEKAL values, respectively.

the two power-laws are given by

\[
\begin{align*}
\bar{T} &\approx 11.6P_{\text{rot}}^{-0.48\pm0.07} \quad \text{MK (MEKAL)}, \\
\bar{T} &\approx 12.2P_{\text{rot}}^{-0.50\pm0.08} \quad \text{MK (APEC)},
\end{align*}
\]

(2.8) \quad (2.9)

where \(P_{\text{rot}}\) is the rotation period in days. Similar results were obtained for the ROSAT data by Güdel et al. (1997a), again considering the higher fit-temperature instead of the mean temperature. As we averaged the temperature with the EM in each bin used as weights, the steady decrease of the mean temperature with period is consistent with a decrease of the amount of hot plasma as the star spins down.

Equations 2.6–2.9 allow us to check for consistency with published relations between \(L_X\) and \(P_{\text{rot}}\). By combining formula (6) with (8) or, respectively, (7) with (9), we find

\[
\begin{align*}
L_X &= 4.01 \times 10^{30}P_{\text{rot}}^{-2.04\pm0.36} \quad \text{erg s}^{-1} \quad \text{(MEKAL)}, \\
L_X &= 4.04 \times 10^{30}P_{\text{rot}}^{-2.03\pm0.35} \quad \text{erg s}^{-1} \quad \text{(APEC)},
\end{align*}
\]

(2.10) \quad (2.11)

which is consistent with previously reported dependences of this type (Pallavicini et al. 1981; Güdel et al. 1997a). A linear regression for \(\log L_X\) and \(\log P_{\text{rot}}\) yields the same result, with a power-law index of \(-2.03\).
2.7. Discussion

We also studied the relation between radio luminosity and the temperature. Radio luminosities or upper limits thereof are available (Güdel & Gaidos 2001b) for five out of the six targets and are plotted in Figure 2.12. They refer to low emission levels outside obvious flares. We use the values of the upper limits in the regression analysis. The slopes of the power-laws are therefore lower limits, given by

\[
L_R \approx 0.86 \times 10^9 T^{5.65 \pm 0.46} \text{ erg s}^{-1} \text{ Hz}^{-1} \text{ (MEKAL)}, \tag{2.12}
\]
\[
L_R \approx 1.69 \times 10^9 T^{5.29 \pm 0.74} \text{ erg s}^{-1} \text{ Hz}^{-1} \text{ (APEC)}. \tag{2.13}
\]

These relations suggest a relation between the non-thermal electron population, responsible for radio gyrosynchrotron emission, and coronal heating.

2.7.2 Abundances

In the solar corona, the so-called FIP effect has been observed, in which the elements with a FIP lower than 10 eV are overabundant relative to the solar photospheric composition, whereas the elements with a higher FIP show the same abundance as the solar photosphere (Feldman 1992; Laming et al. 1995;
Chapter 2. High-Resolution X-Ray Spectroscopy of Solar Analogs

Feldman & Laming 2000). Recent spectroscopic analysis with XMM-Newton and Chandra has shown that in very active stars, an inverse effect is present, in which the low-FIP elements are depleted relative to the high-FIP elements (Brinkman et al. 2001). In our sample, we observe an evolutionary trend from an inverse FIP effect for the most active star 47 Cas B to a solar-like FIP effect in the oldest stars (Figure 2.8). We note, however, that the absolute abundances of low-FIP elements such as Fe do not reach values as high as in the solar corona, where overabundances by factors of a few are common (Feldman 1992).

In Figure 2.13, the abundances of Fe, Ne, and the ratios $A(\text{Ne})/A(\text{Fe})$, $A(\text{O})/A(\text{Ne})$, $A(\text{O})/A(\text{Fe})$, and $A(\text{Mg})/A(\text{Fe})$ are plotted as a function of the temperature (based on MEKAL/SPEX, method 2). The dotted regions include the ranges of a larger stellar sample (Güdel 2004a). The abundance of the low-FIP element Fe tends to decrease from a nearly photospheric value for stars with an average temperature of 3 to 5 MK to a lower abundance of $0.5$ for the two more active stars EK Dra and 47 Cas B. The error bars for the coolest star ($\beta$ Com) are not plotted for Fe and Ne, since they exceed the range illustrated in the figures. This is due to the near-absence of a continuum in this star, which makes absolute abundance determinations difficult. The abundance ratios, however, are robust (see also Audard et al. 2004).

In the middle left and the bottom left panels, the abundance ratio of Ne/Fe and O/Fe, respectively, are shown as a function of the average coronal temperature. Because Ne and O are high-FIP elements, and Fe is a low-FIP element, the ratios increase with temperature, confirming the evolutionary trend of decreasing low-FIP elements (such as Fe) with increasing activity. A similar effect was found by Audard et al. (2003a) for a sample of active RS CVn-type binaries. This is also confirmed by a larger sample of active stars (Güdel 2004a).

The plots in the middle right and bottom right panels show the abundance ratio of O/Ne, two high-FIP elements, and of Mg/Fe, two low-FIP elements. Our sample of stars is too small to constrain a trend in these plots. However, the larger star sample studied by Güdel (2004a) shows a nearly flat distribution for both ratios.

Could it be that the coronal abundance pattern reflects the composition of the underlying photosphere? This view does not find support from other studies. Although a full picture would include knowledge of the abundances of other elements such as C, N, O, we see, from the summary in Sect. 2.2.2, no support for a photospheric abundance pattern that significantly deviates from solar. Finally, we recall that the solar coronal composition does not reflect the photospheric composition either, hence an agreement between photospheric and coronal abundances is not a priori anticipated for solar analogs.
2.7. Discussion

Figure 2.13: Abundances of Fe, Ne, and abundance ratios of Ne/Fe, O/Ne, O/Fe, and Mg/Fe are plotted as a function of the mean coronal temperature. Values from method 2 (SPEX) have been used. The dotted contours delimit the regions derived from a larger stellar sample (Güdel 2004a). Note that the displayed range of the abundances or abundance ratios is 1 dex in each of the six plots.

2.7.3 Flares and Coronal Heating

In the previous sections, we have highlighted correlations between observable parameters, and we have found continuous variability in all six targets to an extent that hardly any time interval is free of fluctuations. Although conventional interpretation of coronal structure often makes use of the approximation
of static coronal loops (as, e.g., described by Rosner et al. 1978), the interpretation of the phenomenology revealed by our light curves cannot be rooted in strictly static loops, although static loop models may, under certain circumstances, serve as approximations even under flaring conditions (Jakimiec et al. 1992).

There is appeal in the alternative, extreme model assuming that the coronal emission is entirely due to dynamic, flaring loops. A number of observed features reported in this work and in the previous literature seem to support such a model: i) with the sensitivity available with XMM-Newton and Chandra, X-ray emission previously ascribed to a quiescent component is now recognized to be continuously variable; in the most extreme cases, no steady component can reasonably be identified in the light curves (Audard et al. 2003b). ii) More active stars (i.e., stars with a higher $L_X/L_{bol}$ ratio) appear to maintain hotter coronae. This is difficult to explain with a model that assumes a corona composed only of steady magnetic loops, with the principal determinant of $L_X$ being the magnetic filling factor, up to the empirical saturation value of $\log L_X/L_{bol} \approx -3$. Such models do not automatically explain why the coronal temperature increases with increasing $L_X$. We discuss this point further in our conclusions below. iii) Active stars continuously produce radio emission from accelerated electrons. The lifetime of the latter is short, probably amounting to no more than seconds to minutes (Kundu et al. 1987). In the solar corona, flare energy release processes are required to produce such electron populations. In order to generate the observed stellar radio emission, high-energy electrons must be replenished frequently.

This question has previously been addressed by studying, for a model in which the corona is heated entirely by flares, the EMD (Güdel 1997c; Güdel et al. 2003), the average temperature to be ascribed to such a corona (Audard et al. 2000), and the light curve characteristics expected from a superposition of stochastic flares (Audard et al. 1999, 2000; Kashyap et al. 2002; Güdel et al. 2003; Arzner & Güdel 2004). We are now in a position to apply the methodology to our results.

First, we use the shapes of our EMDs to characterize the underlying flare population in the framework of this model. Güdel et al. (2003) derived an analytic expression for a DEM of a flare-heated corona under the assumption that the temperature and the flare density both decay exponentially with time constants $\tau_T$ and $\tau_n$, respectively, and that a relation between flare peak-EM and peak-$T$ holds, $EM_p \propto T_p^b$, as reported by Feldman et al. (1995). Then, the DEM follows power-law relations on both sides of its peak temperature ($T_m$), with

$$Q(T) \propto \begin{cases} T^{2/\zeta} & , T \leq T_m \\ T^{-(b-\phi)(\alpha-2\beta)/(1-\beta)+2b-\phi} & , T \geq T_m \end{cases}$$

(2.14)

where $b \approx 4.3 \pm 0.35$ as derived from a large sample of stellar flares (Güdel
2.7. Discussion

The temperature $T_m$ depends on the energy of the smallest flares participating in the heating in this simple model (Güdel et al. 2003). Further, $\beta$ is a power-law index for a relation between the flare e-folding decay time and its radiated energy, $\tau \propto E^\beta$. As noted by Güdel et al. (2003), $\beta$ is probably close to zero although an extreme case of $\beta = 0.25$ was also studied. The variable $\phi$ gives the slope of the cooling function (radiative power per unit emission measure) for a power-law approximation in the temperature interval of interest. We use $\phi = -0.3$ as an approximation of the slope of the cooling function in the temperature interval 6.8–7.5, i.e., the temperature interval above the EMD peak temperature (see Fig. 10 in Audard et al. 2004). The parameter $\alpha$ is of primary interest for us: It is the exponent of the distribution of the occurrence rate $N$ of flares in radiated energy, viz., $dN/dE \propto E^{-\alpha}$ as found for a large sample of solar flares (e.g., Crosby et al. 1993), but applicable also to stellar coronae (see references above). In the first equation, applicable to the cooler portion of the DEM, $\zeta = \tau_n/\tau_T$. This parameter describes the amount of heating occurring during the flare decay, with $\zeta = 2$ corresponding to free cooling without heating, and $\zeta \approx 0.5$ corresponding to extreme heating rates during the decay (Reale et al. 1997).

We have measured the slopes of our EMDs on both sides of $T_m$. To find the possible ranges for the best-fit slopes, we re-analyzed our data using method 2 by starting with different initial conditions for the slopes in our iteration, but the DEMs converged to similar values. The ranges of the resulting best fit slopes (not considering the error ranges in the DEM) are reported in Table 2.10. We then used Eq. 2.14 to determine the most likely $\zeta$ and $\alpha$ with their acceptable ranges. The results are also reported in Table 2.10.

From the low-$T$ slopes of the EMDs derived with method 2 using SPEX, we find $\zeta$ to be around unity in all cases. Such values are typical for individual flares observed on active stars (see Güdel et al. 2003 and references therein) and support our assumption that such flares contribute significantly to the overall observed emission. Static loops, on the other hand, normally produce shallower DEMs, with slopes of $+1$ to $+3/2$, depending on the amount of conductive flux at the loop footpoints (Rosner et al. 1978; van den Oord et al. 1997).

From the high-$T$ slope, we derive $\alpha \approx 2.2 – 2.8$, in excellent agreement with $\alpha$ values determined from long light curves of active stars (Audard et al. 2000; Kashyap et al. 2002; Güdel et al. 2003; Arzner & Güdel 2004). If $\alpha > 2$ and the emitted energy is integrated over the flare-rate distribution to obtain the total radiative loss, i.e., $L_{\text{tot}} = \int_{E_1}^{E_2} E (dN/dE) dE$, then $L_{\text{tot}}$ diverges as $E_1 \to 0$, i.e., the smallest flares dominate coronal heating, and a lower energy cutoff is required for this power-law.

We now continue this consideration by simulating light curves based on a characteristic flare shape for the three most active stars, assuming a flare-rate
Table 2.10: Light curve modeling

<table>
<thead>
<tr>
<th></th>
<th>47 Cas B</th>
<th>EK Dra</th>
<th>$\pi^1$ UMa</th>
</tr>
</thead>
<tbody>
<tr>
<td>Slope for $T &lt; T^a_m$</td>
<td>2.10-2.63</td>
<td>2.22-3.09</td>
<td>1.51-2.46</td>
</tr>
<tr>
<td>Slope for $T &gt; T^b_m$</td>
<td>(-1.34)-(-1.48)</td>
<td>(-1.55)-(-1.62)</td>
<td>(-2.45)-(-3.63)</td>
</tr>
<tr>
<td>$\phi$</td>
<td>-0.3</td>
<td>-0.3</td>
<td>-0.3</td>
</tr>
<tr>
<td>$\zeta$</td>
<td>0.95-0.76</td>
<td>0.90-0.65</td>
<td>1.32-1.23</td>
</tr>
<tr>
<td>$\alpha(\beta = 0)$</td>
<td>2.21-2.28</td>
<td>2.25-2.31</td>
<td>2.43-2.78</td>
</tr>
<tr>
<td>$\alpha(\beta = 0.25)$</td>
<td>2.16-2.20</td>
<td>2.19-2.23</td>
<td>2.33-2.59</td>
</tr>
<tr>
<td>log $E_2/E_1(\beta = 0)$</td>
<td>(-4.2)-(-5.3)</td>
<td>(-2.8)-(-3.4)</td>
<td>(-1.8)-(-2.2)</td>
</tr>
<tr>
<td>log $E_2/E_1(\beta = 0.25)$</td>
<td>(-6.8)-(-6.2)</td>
<td>(-3.5)-(-4.1)</td>
<td>(-1.8)-(-2.5)</td>
</tr>
<tr>
<td>log $E_2$</td>
<td>30.20</td>
<td>30.00</td>
<td>29.0</td>
</tr>
<tr>
<td>mod. depth$^c$</td>
<td>0.028</td>
<td>0.056</td>
<td>0.122</td>
</tr>
<tr>
<td></td>
<td>$\chi^1$ Ori</td>
<td>$\kappa^1$ Cet</td>
<td>$\beta$ Com</td>
</tr>
<tr>
<td>Slope for $T &lt; T^a_m$</td>
<td>1.88-2.86</td>
<td>1.85-2.80</td>
<td>2.12-3.67</td>
</tr>
<tr>
<td>Slope for $T &gt; T^b_m$</td>
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<td>(-2.25)-(-2.60)</td>
<td>(-2.95)-(-3.53)</td>
</tr>
<tr>
<td>$\phi$</td>
<td>-0.3</td>
<td>-0.3</td>
<td>-0.3</td>
</tr>
<tr>
<td>$\zeta$</td>
<td>1.06-0.70</td>
<td>1.08-0.71</td>
<td>0.94-0.54</td>
</tr>
<tr>
<td>$\alpha(\beta = 0)$</td>
<td>2.60-2.80</td>
<td>2.39-2.54</td>
<td>2.54-2.76</td>
</tr>
<tr>
<td>$\alpha(\beta = 0.25)$</td>
<td>2.45-2.60</td>
<td>2.30-2.41</td>
<td>2.40-2.57</td>
</tr>
<tr>
<td>log $E_2/E_1(\beta = 0)$</td>
<td>-</td>
<td>-</td>
<td>-</td>
</tr>
<tr>
<td>log $E_2/E_1(\beta = 0.25)$</td>
<td>-</td>
<td>-</td>
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</tr>
<tr>
<td>log $E_2$</td>
<td>-</td>
<td>-</td>
<td>-</td>
</tr>
<tr>
<td>mod. depth$^c$</td>
<td>-</td>
<td>-</td>
<td>-</td>
</tr>
</tbody>
</table>

$^a$ Determined in the logarithmic energy range between log $T = 6.2$ and (log $T_m$-0.1)

$^b$ Determined in the logarithmic energy range between (log $T_m$+0.1) and an upper limit that depends on the star (log $T_m = 7.5$ for 47 Cas and EK Dra, log $T_m = 7.3$ for $\pi^1$ UMa, and log $T = 7.1$ for the other stars)

$^c$ Modulation depth of the observed light curves

distribution based on $\alpha$ as determined above. The flare shape was derived from the convolution of an exponential function (cut off at $t < 0$, describing the decay) and a Gaussian (important to describe the rise and the peak phase). The shapes of the largest flares in the light curves were used to estimate the characteristic rise and the decay time parameters. We performed two sets of simulations: one with $\beta = 0$ and one with $\beta = 0.25$. In each simulation $\alpha$ was chosen within the acceptable range for a given $\beta$. The flare decay-time was varied according to the scaling $\tau \propto E^\beta$. For the largest flare in each light curve, we measured the amplitude and fitted the decay phase with an exponential function to find the decay time, and hence the emitted energy. We then set the maximum flare energy $E_2$ equal to the energy of the largest observed
2.8 Conclusions

As the rotation rate of a solar analog decreases during its evolution on the main sequence, the efficiency of the internal dynamo weakens, resulting in a decrease of the magnetic activity in the stellar atmosphere. We have studied systematic trends in the long-term evolution of stellar coronal X-ray emission for ages in the range 0.1–2 Gyr. As a consequence of the stellar spin-down,
Figure 2.14: Simulated light curves obtained assuming a flare distribution that is based on the α values found from our EMDs. The maximum flare energy and the difference between largest and smallest flares assumed here are given in Table 2.10. The black shapes represent the largest flares actually used to synthesize the light curves from the simulations (and this flare may therefore be smaller than the largest flare actually observed in the light curves in Fig. 2.9, which are not considered for this comparison).

the X-ray luminosity steadily decreases from levels that may be close to the empirical saturation limit in the youngest stars \( L_X / L_{bol} \approx 10^{-3} \), to levels approximately two orders of magnitude lower within the first two Gyr. During the next \( \approx 3-4 \) Gyr, \( L_X \) reduces by another factor of ten to levels as seen in α Cen, β Hyi, or the Sun (Güdel et al. 1997a). The overall trends reported here confirm earlier studies based on lower-resolution X-ray spectroscopy (Maggio et al. 1987; Güdel et al. 1997a).

The high-resolution X-ray spectroscopy now available has permitted a more detailed study of the composition and the thermal structure of solar-like coronae than was previously possible. We have studied emission measure distributions and coronal element abundances for all six targets in a homogeneous way, applying two widely differing methods and using two different sets of atomic parameters. It is important to recall that both of our methods include modeling of line blends to the extent possible with the presently available line emissivities.
There is gratifying agreement between the results from the two methods, although it appears that the choice of the atomic database introduces systematic differences. This is perhaps not surprising as the presently available compilations of atomic parameters are incomplete and suffer from systematic uncertainties. A more serious limitation is set by the mathematical problem of spectral inversion itself. While even counting statistics at the percent level makes the inversion problem ill-conditioned (Craig & Brown 1976; Judge 2002), systematic uncertainties in the line emissivities of perhaps up to 10–20% may introduce various structure in our EMDs that may not correspond to coronal features. We have suggested that a reasonable convergence criterion should be set, although it is difficult to assess at what level artificial structure is introduced into a reconstructed EMD. Regardless of these inconsistencies, however, we recover element abundances that are rather robust.

With these limitations in mind, we have found systematic trends in the EMD structure as a star ages. Not only does the total emission measure continuously decrease, the temperature where the EMD peaks also decays with time. The EM-weighted logarithmic average of the coronal temperature $\bar{T}$ thus follows a power-law dependence on the X-ray luminosity, namely $\bar{T} \propto L_X^{0.25\pm0.02}$. A similar relation between the dominant coronal temperature and $L_X$ was already studied by Schrijver et al. (1984), using Einstein data of a large sample of stars. Schmitt et al. (1990) found for 1-T model fits $T \propto L_X^{0.4}$ from Einstein data. Finally, Güdel et al. (1997a) used ROSAT data and a G-star sample similar to ours and found a relation between the higher temperature of a 2-T model and $L_X$ in complete agreement with our results.

The observed trends cannot be explained by a model that is based exclusively on different filling factors of the surface magnetic field, as an increased filling factor does not explain per se why the temperature should increase. We have discussed an extreme case of an alternative model in which a statistical distribution of \textit{flare} is responsible for the correlation between $L_X$ and $\bar{T}$. The question then is why the flare rate (above any given base-level energy) is higher in more active (i.e., more rapidly rotating) stars. Our data cannot give a conclusive answer. One possibility is that more active stars indeed do show a larger magnetic surface filling factor, and that the higher density of magnetic loops leads to more magnetic reconnection, thus producing a higher flare rate (Güdel et al. 1997a); this, then, also includes a more prominent population of very large flares that produce both large emission measures and very high temperatures, thus shifting the average temperature to higher values. Once the magnetic filling dilutes, the interactions between neighboring magnetic loop systems will become less frequent, and both $L_X$ and $\bar{T}$ decrease (see Güdel et al. 1997a).

Alternatively, instead of increasing the surface filling factor, other ingredients may systematically change with changing rotation period, such as the structure of surface magnetic field in active regions or differences in the con-
vection pattern that jostles the magnetic-loop footpoints. In a statistical-flare model, then, the way to explain the $L_X - \theta^T$ correlation would be to reheat the same active regions more frequently in more active stars, as some process brings non-potential energy into the magnetic fields at a higher rate. As argued by Audard et al. (2000), such frequently-heated loops can be approximated by static loops although the heating process is non-static. In this case, approximating the loop temperature with its apex temperature (where most of the EM is found), the Rosner et al. (1978) loop scaling law predicts $L_X \propto T^{3.7}$ for loops of given length below the coronal pressure scale height, and $L_X \propto T^{4.7}$ for loops larger than this limit (see Audard et al. 2000 for details). These predictions are close to our observational finding. On the other hand, static loops would predict slopes of the DEM on the low-temperature side that are significantly smaller than those determined by us.

To conclude, we are presently unable to distinguish between a model in which the flare rate is controlled by the magnetic filling factor, and one in which constrained active regions flare progressively more frequently as the rotation rate of a star increases. Both approaches, however, are compatible with the hypothesis that much of the coronal heating is induced by flaring, regardless of the ultimate cause of the increased flare rate in more active stars.

Because larger flares produce hotter plasma (Feldman et al. 1995), more active stars produce hotter coronae. This trend is unequivocally recovered from our observations and further supports a picture in which flares contribute significantly to the overall coronal heating (e.g., Audard et al. 2000; Kashyap et al. 2002; Güdel et al. 2003).

We note in passing that an alternative view with similar consequences has recently been presented by Peres et al. (2004). These authors suggest that, based on properties of coronal structures seen on the Sun, a higher occurrence of very compact, hot features including flares make more active coronae hotter.

We have also derived element abundances and found good agreement between our two methods. Somewhat more systematic deviations can be noted if different atomic databases are used (MEKAL, APEC), but the trends in the abundance pattern agree and markedly change with changing activity level. A similar trend was noted for RS CVn binaries by Audard et al. (2003a), but the latter study referred to extremely active stars in which the abundance pattern changed from a strong inverse FIP effect to a flat distribution with decreasing activity (or mean coronal temperature). In our sample, the change from an inverse or flat distribution to a solar-like distribution occurs at ages of less than 300 Myr or rotation periods longer than $\approx 3$ days. Incidentally, a rapid decay of non-thermal radio emission has been noted for the same activity range. We hypothesize that the same electrons that are responsible for the observed gyrosynchrotron emission also induce an inverse-FIP effect in the most active stars, as follows (see Güdel et al. 2002 for further arguments): if electrons are streaming along the magnetic fields toward the chromosphere,
they build up a downward-pointing electric field that acts to suppress positive currents from the chromosphere to the corona. In other words, ions in the chromosphere are prevented from streaming into the corona, while neutral, predominantly high-FIP elements, are not affected. As the electron population diminishes in less active stars, the suppression of ion diffusion into the corona disappears, and a solar-like FIP effect can build up, by whatever (still unidentified) mechanisms. Recently, Laming (2004) presented an alternative model in which both the solar-like FIP and the inverse FIP effect are related to a common plasma-physical cause.
Chapter 2. *High-Resolution X-Ray Spectroscopy of Solar Analogs*
Chapter 3

High-Resolution X-Ray Spectroscopy of T Tauri Stars in the Taurus-Auriga Complex

**ABSTRACT:** Differences have been reported between the X-ray emission of accreting and non-accreting stars. Some observations have suggested that accretion shocks could be responsible for part of the X-ray emission in classical T Tauri stars (CTTS). We present high-resolution X-ray spectroscopy of nine pre-main sequence stars in order to test the proposed spectroscopic differences between accreting and non-accreting pre-main sequence stars. We use X-ray spectroscopy from the *XMM-Newton* Reflection Grating Spectrometers and the EPIC instruments. We interpret the spectra using optically thin thermal models with variable abundances, together with an absorption column density. For BP Tau and AB Aur we derive electron densities from the O\textsuperscript{vii} triplets. Using the O\textsuperscript{vii}/O\textsuperscript{viii} count ratios as a diagnostic for cool plasma, we find that CTTS display a soft excess (with equivalent electron temperatures of $\approx 2.5 - 3$ MK) when compared with WTTS or zero-age main-sequence stars. Although the O\textsuperscript{vii} triplet in BP Tau is consistent with a high electron density ($3.4 \times 10^{11}$ cm$^{-3}$), we find a low density for the accreting Herbig star AB Aur ($n_e < 10^{10}$ cm$^{-3}$). The element abundances of accreting and non-accreting stars are similar. The Ne abundance is found to be high (4-6 times the Fe abundance) in all K and M-type stars. In contrast, for the three G-type stars (SU Aur, HD 283572, and HP Tau/G2), we find an enhanced Fe abundance (0.4-0.8 times solar photospheric values) compared to later-type stars. Adding the results from our sample to former high-resolution studies of T Tauri stars, we find a soft excess in all accreting stars, but in none of the non-accretors. On the other hand, high electron densities and high Ne/Fe abundance ratios do not seem to be present in all accreting pre-main sequence stars.
3.1 Introduction

Young, late-type stars are known to be luminous X-ray emitters. This is particularly true for T Tauri stars, which are divided into two classes: the classical T Tauri stars (CTTS) and the weak-line T Tauri stars (WTTS). The CTTS display strong H\textalpha emission lines as a signature of accretion and an infrared excess as a signature of the presence of a circumstellar disk. The WTTS display much weaker H\textalpha lines and no infrared excess, a sign that accretion has ceased and a thick disk is no longer present. While CTTS are thought to be in an earlier evolutionary stage, both types of stars are simultaneously present in star forming regions.

X-rays from T Tauri stars were first detected in early studies with the Einstein satellite (Feigelson & DeCampli 1981) and surveyed more extensively with the ROSAT satellite (e.g., Feigelson et al. 1993; Neuhauser et al. 1995). Both types of young stars exhibit variability on time scales of order of hours to days. Their overall X-ray properties were found to be similar to X-ray characteristics of main-sequence stars, and the X-ray emission was interpreted to arise from a corona.

However, most young stellar objects, especially in their early evolutionary stage, are believed to be fully convective, therefore the generation of magnetic fields through a solar-like dynamo should not be possible. Also, the presence of a disk and accretion in CTTS adds another important element: shocks generated by accretion of circumstellar material onto the star could contribute to the X-ray emission.

The influence of accretion on X-ray emission processes has been debated. In the ROSAT observations of the Taurus-Auriga complex, Stelzer & Neuhauser (2001) found that the $L_X/L_*$ ratio (where $L_*$ is the bolometric luminosity of the star) is lower in CTTS than in WTTS. This has been confirmed in the recent, more comprehensive XMM-Newton Extended Survey of the Taurus Molecular Cloud (XEST), but on the other hand, no direct correlation is found between accretion rate and $L_X/L_*$ (see Chapter 4). Similar results have also been reported from the Chandra Orion Ultradeep Project (COUP), where the CTTS are found to be less luminous in X-rays than WTTS (Preibisch et al. 2005). In both surveys, the bulk X-ray emission is consistent with coronal emission.

On the other hand, recent results from high-resolution X-ray spectroscopy of some CTTS seem to indicate the presence of X-ray emission generated in accretion shocks. The most evident case is TW Hya. The Chandra and XMM-Newton spectra of TW Hya (Kastner et al. 2002; Stelzer & Schmitt 2004) are dominated by emission from plasma at temperatures of $\approx 3$ MK which is much lower than usually found in T Tauri stars. The O\textsc{vii} triplet, consisting of lines at 21.6 \AA\ (resonance, $r$ line), 21.8 \AA\ (intercombination, $i$ line), and 22.1
3.1. Introduction

(forbidden, $f$ line), displays an $f/i$ flux ratio much below unity. This triplet is density-sensitive (Gabriel & Jordan 1969); in late type stars, typical values for the $f/i$ ratio are higher than unity (Ness et al. 2004; Testa et al. 2004), indicating densities of at most a few times $10^{10}$ cm$^{-3}$. In TW Hya, the measured electron density exceeds $10^{12}$ cm$^{-3}$ (Kastner et al. 2002; Stelzer & Schmitt 2004), i.e., it is at least two orders of magnitude higher than typical coronal densities. Studies of other density-sensitive lines lead to similar results: the He-like Ne IX triplet in the Chandra observations gives log $n_e = 12.75$ (Kastner et al. 2002), while a study by Ness & Schmitt (2005) presents additional evidence of high density from the flux ratios of Fe XVII lines. Further, Kastner et al. (2002) and Stelzer & Schmitt (2004) found an abundance anomaly in the spectrum of TW Hya, with certain metals being underabundant with respect to the solar photospheric abundances, while nitrogen and neon are found to be strongly overabundant. This anomaly was interpreted by Stelzer & Schmitt (2004) as being due either to metal depletion by condensation onto grains in the accretion disk or to an abundance anomaly present in the original molecular cloud. A particularly high Ne/Fe abundance ratio was also found by Argiroffi et al. (2005) for TWA 5, a stellar system in the TW Hya association. This object is a quadruple system and one of the components could be an accreting star (Mohanty et al. 2003), although it is not possible to identify which of the stars is the X-ray source. The high Ne/Fe ratio could be due to the accreting star, but could also be an environmental effect of the TW Hydrae association (Argiroffi et al. 2005). Drake et al. (2005b) have proposed to use the Ne/O abundance ratio as a diagnostic for metal depletion in accreting stars. This ratio is in fact found to be substantially higher in TW Hya than in all other studied stars (also higher than in TWA 5, another star of the TW Hydrae association).

Schmitt et al. (2005) have discussed the high-resolution X-ray spectrum of the CTTS BP Tau observed with XMM-Newton. This spectrum also displays an $f/i$ ratio lower than unity, resulting in an electron density of log $n_e \approx 11.5$. However, the BP Tau spectrum is dominated by a hot plasma component. The authors interpreted the hotter component as originating from a corona or from magnetic interaction between disk and corona, while the soft component may originate from accretion shocks.

Robrade & Schmitt (2006) have presented a comparative study of high-resolution spectra of four pre-main sequence accreting stars: BP Tau, CR Cha, SU Aur, and TW Hya. They tentatively added CR Cha as another CTTS with a low O VII $f/i$ ratio, but the low signal-to-noise (S/N) ratio makes the measurements uncertain. The O VII triplet was not detected in SU Aur. The Ne abundance is found to be enhanced relative to Fe also in CR Cha and BP Tau. However, the Ne/O abundance ratio of BP Tau is similar to the ratio found for other (non-accreting) stars and not as high as that of TW Hya, which
may be due to evolutionary effects in the accretion disk (Drake et al. 2005b).

Recently, Günther et al. (2006) studied the Chandra spectrum of another CTTS, V4046 Sgr, and measured a high O\textsc{vii} f/i ratio, consistent with electron densities of $\log n_e \approx 11.5$.

High-resolution spectroscopy is of crucial importance to identify the soft component in X-ray spectra of young pre-main sequence stars because emission lines formed at low coronal temperatures can be accessed, such as lines of C\textsc{v}, C\textsc{vi}, N\textsc{vi}, N\textsc{vii}, O\textsc{vii}, and O\textsc{viii}, whereas the energy resolution of CCDs is insufficient to reveal these lines individually. The possibility to measure density-sensitive lines is also essential to distinguish between X-ray emission that may originate in accretion shocks and coronal emission. A comparative study of X-ray spectra of CTTS and WTTS should further our understanding of the role of accretion in the production of X-ray emission.

In this chapter we present high-resolution X-ray spectra of four CTTS, four WTTS, and a one Herbig Ae/Be star. These observations are part of the XEST project (Güdel et al. 2007a), the XMM-\textit{Newton} survey of the most populated regions in the Taurus Molecular Cloud. The purpose of that survey is to study X-ray emission in a large fraction of the TMC population. While the spectra of HD 283572, SU Aur, and BP Tau have been extracted from archival data and have already been presented in the literature (Scelsi et al. 2005; Schmitt et al. 2005; Robrade & Schmitt 2006), the RGS spectra of the other stars are shown for the first time in this chapter. We reanalyze the spectra of the former three stars to provide a consistent comparison.

The structure of this chapter is as follows. The stellar sample is described in Section 3.2, while we present our observations and data reduction in Sect. 3.3. The results are presented in Sect. 3.4, while a detailed discussion of the O\textsc{vii} triplets is presented in Sect. 3.5. In Sect. 3.6 we discuss our results and Sect. 3.7 contains our conclusions.

### 3.2 Stellar Sample

We present RGS spectra of 9 pre-main sequence stars. Four of them are CTTS, four are WTTS, while one is an intermediate-mass young star, belonging to the class of Herbig stars (Herbig 1960). The principal properties of the stars are listed in Table 3.1. We discuss some specific characteristics below.
Table 3.1: Properties of target stars. $L_\ast$ is the stellar (photospheric) bolometric luminosity, $M$ is the mass, $A_V$ and $A_J$ are the extinctions in the $V$ and $J$ bands, respectively, $P_{\text{rot}}$ is the rotation period, $\text{EW}(\text{H}\alpha)$ is the equivalent width of the H\alpha line (positive for emission), and $\dot{M}$ is the mass accretion rate. If not otherwise noted, the properties are reported from Güdel et al. (2007a) and references therein.

<table>
<thead>
<tr>
<th>Star</th>
<th>XEST No.$^a$</th>
<th>$L_\ast$ ($L_\odot$)</th>
<th>$M$ ($M_\odot$)</th>
<th>Age (Myr)</th>
<th>$R$ ($R_\odot$)</th>
<th>$A_V$ (mag)</th>
<th>$A_J$ (mag)</th>
<th>$P_{\text{rot}}$ (d)</th>
<th>Spec.</th>
<th>$\text{EW}(\text{H}\alpha)$ (Å)</th>
<th>$\log \dot{M}$ ($M_\odot , \text{yr}^{-1}$)</th>
<th>Type</th>
<th>TTS</th>
</tr>
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<tbody>
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<td>HD 283572</td>
<td>21-039</td>
<td>6.50</td>
<td>1.70</td>
<td>7.92</td>
<td>2.56</td>
<td>0.38</td>
<td>0.11</td>
<td>1.55</td>
<td>G5</td>
<td>0</td>
<td>–</td>
<td>W</td>
<td></td>
</tr>
<tr>
<td>V773 Tau</td>
<td>20-042</td>
<td>5.60</td>
<td>1.53</td>
<td>6.35</td>
<td>1.91</td>
<td>1.39</td>
<td>0.31</td>
<td>3.43</td>
<td>K2</td>
<td>4-10</td>
<td>$&lt;-10$</td>
<td>W</td>
<td></td>
</tr>
<tr>
<td>V410 Tau$^c$</td>
<td>23-032/24-028</td>
<td>2.20</td>
<td>1.51</td>
<td>2.74</td>
<td>2.31</td>
<td>0.67</td>
<td>0.00</td>
<td>1.94</td>
<td>K4</td>
<td>2-3</td>
<td>$&lt;-8.8$</td>
<td>W</td>
<td></td>
</tr>
<tr>
<td>HP Tau/G2</td>
<td>08-051</td>
<td>6.50</td>
<td>1.58</td>
<td>10.5</td>
<td>2.34</td>
<td>2.08</td>
<td>0.66</td>
<td>1.20</td>
<td>G0</td>
<td>0-5</td>
<td>–</td>
<td>W</td>
<td></td>
</tr>
<tr>
<td>SU Aur</td>
<td>26-067</td>
<td>9.90</td>
<td>1.91</td>
<td>6.02</td>
<td>3.06</td>
<td>0.90</td>
<td>0.21</td>
<td>1.70</td>
<td>G2</td>
<td>2-6</td>
<td>-8.30/-8.20</td>
<td>C</td>
<td></td>
</tr>
<tr>
<td>DH Tau</td>
<td>15-040</td>
<td>0.56</td>
<td>0.47</td>
<td>1.53</td>
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<td>0.32</td>
<td>7.00</td>
<td>M1</td>
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<td>C</td>
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<td>BP Tau</td>
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<td>K7</td>
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</tr>
<tr>
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<td>$&lt;1.46$</td>
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<td>22-44</td>
<td>$\approx -8^b$</td>
<td>Ae</td>
<td></td>
</tr>
</tbody>
</table>

$^a$ XEST catalog number: the first two digits define the exposure numbers, the last three digits are the source numbers (Güdel et al. 2007a).

$^b$ From Chapter 5 and references therein.

$^c$ V410 Tau was observed in two different exposures, therefore two different XEST numbers are given (see below).
Chapter 3. High-Resolution X-Ray Spectroscopy of TTS in Taurus-Auriga

HD 283572 is a single G-type star that does not show signs of accretion and is therefore classified as a WTTS (Kenyon et al. 1998). Also, there is no evidence of an infrared excess that would be due to a circumstellar dust disk (Furlan et al. 2006). Favata et al. (1998) have studied its X-ray variability, analyzing data from *Einstein, ASCA, ROSAT*, and SAX. Results from the *XMM-Newton* data set have already been presented by Scelsi et al. (2005).

V773 Tau is a system composed of four stars. The system was detected to be a binary by Ghez et al. (1993) and Leinert et al. (1993); the primary component was then discovered to be a binary itself (Welty 1995). Recently, a fourth component has been detected by Duchêne et al. (2003). These authors showed that the two close components A and B reveal properties of WTTS, while the C component shows a near infrared excess typical of CTTS. The fourth component has been classified as an infrared companion (IRC). The nature of IRC is debated: they could be deeply embedded TTS undergoing strong accretion (see e.g. Koresko et al. 1997) or they may be embedded protostars (e.g. Ressler & Barsony 2001). In the latter case, the four stars would be in three different evolutionary phases. We refer the reader to Duchêne et al. (2003) for an exhaustive discussion. In this work we will treat V773 Tau as a WTTS for two reasons: the primary and secondary stars (the most luminous components) are WTTS, and the hydrogen column density ($N_H$) found in our spectral fits ($2.0 \times 10^{21}$ cm$^{-2}$) is consistent with the optical extinctions found for the two WTTS ($A_V = 1.39$ mag) if a standard gas-to-dust ratio is assumed ($N_H/A_V = 2 \times 10^{21}$ cm$^{-2}$ mag$^{-1}$; Vuong et al. 2003). The masses of V773 Tau A and B are 1.5 and $\approx 1 M_\odot$, respectively, while the mass of the C component is only $\approx 0.7 M_\odot$. The IRC would have a mass of $\leq 0.7 M_\odot$. Because the X-ray luminosities of TTS are correlated with mass (see Preibisch et al. 2005 and Chapter 4), we expect the A+B components to dominate the X-ray spectrum. X-rays from V773 Tau were previously detected with *ROSAT* (Feigelson et al. 1994) and with *ASCA* (Skinner et al. 1997).

V410 Tau is a triple system (Ghez et al. 1993, 1997). V410 Tau A and B are separated by 0.12'' with B being much fainter. The C component is also much fainter than V410 Tau A. The primary is of spectral type K4 with a mass of 1.5 $M_\odot$, and has an age of 2.74 Myr. V410 Tau was detected in X-rays by *ROSAT* (Strom & Strom 1994), and more recently Stelzer et al. (2003) presented a set of *Chandra* observations of this source.

SU Aur is a single star of spectral type G2, with a mass of 1.9 $M_\odot$. The properties of this object are described by DeWarf et al. (2003). Despite the low equivalent width (EW) of the Hα line reported in Table 3.1, its early spectral type and the measured infrared excess classifies SU Aur as a CTTS (Muzerolle et al. 2003). The source is one of the X-ray brightest CTTS and was already detected by Feigelson & DeCampli (1981) with the *Einstein Observatory*, and by Skinner & Walter (1998) in an *ASCA* observation. Results from its *XMM-Newton* spectrum have recently been published by Robrade & Schmitt (2006, 
further analysis will be presented by Franciosini et al. 2007, in preparation).

HP Tau/G2 is a G0 star that forms a triple system with HP Tau/G3 (separation 10")\), the latter itself being a binary (Richichi et al. 1994). The X-ray spectrum of HP Tau/G2 is contaminated by HP Tau/G3. However, fitting separate PSF in the EPIC data, we found that the source counts of HP Tau/G3 amount to only about 8\% of the count rates of HP Tau/G2 (Güdel et al. 2007a). The contamination due to HP Tau/G3 is therefore negligible. Bipolar outflows have been detected from HP Tau/G2 (Duvert et al. 2000), although the star is classified as a WTTS.

DH Tau is another binary, the primary DH Tau A being a CTTS, with a mass of 0.5 $M_\odot$ and an age of 1.5 Myr. DH Tau B is a brown dwarf (Itoh et al. 2005) with mass of 0.03-0.04 $M_\odot$ and an age of 3-10 Myr.

BP Tau is a CTTS of spectral type K7 with a mass of 0.75 $M_\odot$ and an age of $\approx$ 2 Myr, according to the evolutionary model of Siess et al. (2000). The star was already observed by the Einstein Observatory, with $L_X \approx 1 \times 10^{30}$ erg s$^{-1}$ (Walter & Kuhi 1981). Results from the XMM-Newton spectrum have been discussed by Schmitt et al. (2005) and Robrade & Schmitt (2006). They measured a relatively high density in the O\textsc{vii} triplet, suggesting that some of the X-rays could be produced in accretion shocks.

DN Tau is a CTTS of spectral type M0 with a mass of 0.56 $M_\odot$ and an age of $\approx$ 1 Myr. According to Muzerolle et al. (2003), the inclination angle is 35°, and the disk is relatively small (0.05 AU in radius).

AB Aur is a Herbig star of spectral type B9.5-A0 with a mass of 2.7 $M_\odot$ and an age of $\approx$ 4 Myr (DeWarf & Fitzpatrick, private communication). It is surrounded by a disk from which it is accreting material (see for example Catala et al. 1999): it is therefore useful to compare the X-ray emission of this star with the X-ray emission of CTTS. A detailed discussion of the X-ray spectrum of this star is presented in Chapter 5.
Table 3.2: Observation log

<table>
<thead>
<tr>
<th>Stars</th>
<th>XEST No.</th>
<th>ObsID(^1)</th>
<th>Instruments</th>
<th>EPIC Filter</th>
<th>Start Time</th>
<th>Stop Time</th>
<th>Exposure [s]</th>
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<tr>
<td>V773 Tau</td>
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<td>0203542001</td>
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<td>2004-09-12 15:52:37</td>
<td>31674</td>
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<td>0086360301</td>
<td>RGS1, RGS2, MOS1</td>
<td>Medium</td>
<td>2001-03-11 12:46:45</td>
<td>2001-03-12 09:13:24</td>
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<td>0086360401</td>
<td>RGS1, RGS2, MOS1</td>
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<td>2001-03-12 09:29:38</td>
<td>2001-03-12 21:47:51</td>
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<tr>
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<td>2004-08-26 18:10:59</td>
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<tr>
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<td>0101440801</td>
<td>RGS1, RGS2, MOS1</td>
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</tr>
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<td>RGS1, RGS2, MOS1</td>
<td>Thick</td>
<td>2004-08-15 06:14:30</td>
<td>2004-08-16 18:42:57</td>
<td>131307</td>
</tr>
<tr>
<td>DN Tau</td>
<td>12-040</td>
<td>0203542101</td>
<td>RGS1, RGS2, MOS1</td>
<td>Medium</td>
<td>2005-03-04 20:22:29</td>
<td>2005-03-05 05:01:54</td>
<td>31165</td>
</tr>
<tr>
<td>AB Aur</td>
<td>26 043</td>
<td>0101440801</td>
<td>RGS1, RGS2, MOS2</td>
<td>Thick</td>
<td>2001-09-21 01:34:17</td>
<td>2001-09-22 13:34:31</td>
<td>129614</td>
</tr>
</tbody>
</table>

\(^1\) XMM-Newton observation identification number
3.3 Observations and Data Analysis

The XEST campaign collected 28 *XMM-Newton* observations within the Taurus-Auriga complex (Güdel et al. 2007a). We obtained high-resolution X-ray spectra from nine targets within these fields of view, which are in the focus of this chapter. The observation log is given in Table 3.2. For each star, we used each spectrum from the two Reflecting Grating Spectrometers (RGS, den Herder et al. 2001) and one of the spectra from the MOS-type European Photon Imaging Cameras (EPIC, Turner et al. 2001). For HD 283572, the RGS1 instrument was out of operation and, therefore, we restrict our analysis to the RGS2 and MOS1 instruments. For all observations, the MOS instrument observed in full frame mode. Its detectors are sensitive in the energy range of 0.15–15.0 keV with a spectral resolving power of $E/\Delta E = 20 - 50$. The RGSs are suited for high-resolution spectroscopy, operating in the wavelength range of 6–35 Å with a resolution of $\Delta \lambda \approx 60 - 76$ Å.

We reduced our data using the Science Analysis System (SAS) version 6.1. For the EPIC detectors, the data were reduced using the *emchain* task and the sources were detected using the maximum likelihood algorithm *emldetect* (see Güdel et al. 2007a for more details). We reduced the RGS data using the task *rgsproc*. The calibration files used in the data reduction are those described by Pollock (2004). We extracted the total (source+background) spectra and the background spectra separately. In order to optimize the signal-to-noise (S/N) ratio of the spectra and given the weakness of most of them, we decided to include only 85% of the cross-dispersion Point Spread Function (PSF, *xpsfinc=85*). For the background extraction, we kept the exclusion region of the cross-dispersion PSF and the inclusion region of the pulse-height distribution at the default values, namely 95% and 90%, respectively (*xpsfexcl=95 and pdistincl=90*).

We fitted the spectra in XSPEC (Arnaud 1996), using the optically-thin collisionally-ionized plasma model calculated with the Astrophysical Plasma Emission Code (APEC, Smith et al. 2001). In order to account for calibration discrepancies between the RGS and the MOS detectors, we introduced effective area factors fixed at 1.0 for the MOS and 1.05 for the RGS, applicable to our wavelength region according to Kirsch et al. (2004).

Because we are mainly interested in the line-dominated RGS spectra, we used, for the fit procedure, both RGS spectra (if available) but only one MOS spectrum (MOS2 for AB Aur and MOS1 for all other stars), the latter confined to short wavelengths (between 1.5 and 9.35 Å). This range is important to obtain abundances of Mg, Si, S, and Fe. For the RGS spectrum, we used the wavelength region between 8.3 Å and 25.0 Å, except for AB Aur, where we fitted RGS1 between 10 Å and 28 Å and RGS2 between 8.3 Å and 26.5 Å, in order to include the N vii line at 24.78 Å (we choose slightly different
Chapter 3. High-Resolution X-Ray Spectroscopy of TTS in Taurus-Auriga

Wavelength intervals in order to exclude the spectral ranges where an accurate background subtraction was most difficult. A similar approach has been followed in Chapter 2 and by Audard et al. (2003a).

In order to use the \( \chi^2 \) statistic for the fitting procedure, the total spectra were binned to a minimum of 20 counts per bin for RGS (15 counts per bin for the faint spectrum of DN Tau) and to a minimum of 15 counts per bin for MOS.

The spectra were analyzed using two different approaches. First, we fitted the data with a model consisting of a continuous emission measure distribution (EMD) approximated by two power laws as used in the fits to the EPIC spectra in the XEST survey (Güdel et al. 2007a; Chapter 1, Sect. 1.6). The EMD model is approximated by a grid of isothermal components, defining two power laws, one at low temperatures and one at high temperatures. This model was motivated by the EMD shape found from previous high-resolution X-ray spectroscopy of young solar analog stars (see Chapter 2). The model can be described by

\[
Q(T) = \begin{cases} 
EM_0 \cdot (T/T_0)^\alpha & \text{for } T \leq T_0 \\
EM_0 \cdot (T/T_0)^\beta & \text{for } T > T_0
\end{cases}
\]

(3.1)

where \( T_0 \) is the temperature at which the power laws cross, and \( EM_0 \) is the emission measure per log \( T \) at this temperature. The slopes of the power laws below and above \( T_0 \) are \( \alpha \) and \( \beta \), respectively. We left \( \beta \) free to vary (between \( -3 \leq \beta \leq 1 \)), while we fixed the slope \( \alpha \) at a value of 2, as suggested from EMDs derived in previous studies (Chapter 2; Argiroffi et al. 2004). We set low and high temperature cut-offs of the two power laws at \( \log T = 6.0 \) and \( \log T = 8.0 \). The parameters that we fitted hence are \( T_0, EM_0, \beta \) and a selection of elemental abundances.

Although our EMD model is constrained by some approximations (double power-law shape, \( \alpha \) fixed at +2), for a coronal source it is physically more plausible than a 2 or 3-temperature model, as the emission is described by a continuous set of temperatures rather than by only 2 or 3 isothermal components. However, because of the approximations, the EMD model could be inadequate if the X-rays were produced, for example, in near-isothermal accretion shocks. Therefore, we further test our results with a model with two or three isothermal components. We also apply an EMD model in which \( \alpha \) is a free parameter.

In both models, we generally fixed the absorption column density \( N_H \) at the value found in the XEST survey from the EPIC data (Table 3.3). These values agree well with \( A_V \) measurements from the literature (see Table 3.1) if a standard gas-to-dust ratio is assumed. We decided to fit \( N_H \) only for the high-quality spectrum of SU Aur. The value found in the XEST survey in fact led to problems with some RGS lines, which was not the case when we fitted \( N_H \). The \( N_H \) found in the fits of SU Aur agrees much better with the
Table 3.3: Target stars: Results from the EPIC spectral fits. $L_X$ is calculated in the range 0.3-10 keV

<table>
<thead>
<tr>
<th>Stars</th>
<th>$N_H$ $(10^{22} \text{ cm}^{-2})$</th>
<th>$T_0$ (MK)</th>
<th>$\beta$</th>
<th>EM$^1$</th>
<th>$T_{av}$ (MK)</th>
<th>$L_X^2$</th>
</tr>
</thead>
<tbody>
<tr>
<td>HD 283572</td>
<td>0.08 (0.07, 0.08)</td>
<td>10.4 (10.0, 10.7)</td>
<td>-0.94 (-1.02, -0.87)</td>
<td>114.3</td>
<td>14.4</td>
<td>13.0</td>
</tr>
<tr>
<td>V 773 Tau</td>
<td>0.17 (0.17, 0.17)</td>
<td>8.7 (8.4, 9.1)</td>
<td>-0.87 (-0.93, -0.82)</td>
<td>89.8</td>
<td>13.0</td>
<td>9.5</td>
</tr>
<tr>
<td>V 410 Tau</td>
<td>0.02 (0.02, 0.03)</td>
<td>10.4 (9.4, 10.8)</td>
<td>-1.17 (-1.28, -0.96)</td>
<td>44.5</td>
<td>13.2</td>
<td>4.7</td>
</tr>
<tr>
<td>HP Tau/G2</td>
<td>0.41 (0.39, 0.42)</td>
<td>8.0 (8.6, 9.9)</td>
<td>-1.37 (-1.46, -1.28)</td>
<td>146.6</td>
<td>11.0</td>
<td>9.7</td>
</tr>
<tr>
<td>SU Aur</td>
<td>0.47 (0.43, 0.48)</td>
<td>6.4 (6.2, 7.6)</td>
<td>-1.11 (-1.21, -1.06)</td>
<td>95.4</td>
<td>8.9</td>
<td>9.5</td>
</tr>
<tr>
<td>DH Tau</td>
<td>0.20 (0.19, 0.21)</td>
<td>11.5 (11.0, 12.1)</td>
<td>-1.38 (-1.49, -1.26)</td>
<td>80.6</td>
<td>13.4</td>
<td>8.5</td>
</tr>
<tr>
<td>BP Tau</td>
<td>0.06 (0.06, 0.07)</td>
<td>7.1 (6.7, 7.6)</td>
<td>-0.67 (-0.73, -0.60)</td>
<td>12.8</td>
<td>12.3</td>
<td>1.4</td>
</tr>
<tr>
<td>DN Tau</td>
<td>0.07 (0.07, 0.08)</td>
<td>10.5 (9.5, 11.5)</td>
<td>-1.25 (-1.51, -1.02)</td>
<td>11.0</td>
<td>12.9</td>
<td>1.1</td>
</tr>
<tr>
<td>AB Aur</td>
<td>0.05 (0.03, 0.07)</td>
<td>4.8 (4.3, 5.5)</td>
<td>-1.55 (-1.78, -1.41)</td>
<td>3.8</td>
<td>5.7</td>
<td>0.4</td>
</tr>
</tbody>
</table>

1 Total EM integrated over temperature bins between $\log T[K] = 6 - 7.9$ (see Güdel et al. 2007a for more details), given in units of $10^{52} \text{ cm}^{-3}$.

2 Given in units of $10^{30} \text{ erg s}^{-1}$.

$A_V$ values given in the literature ($A_V = 0.9$ mag, Kenyon & Hartmann 1995).

We fitted elemental abundances for the lines that are clearly seen in the spectra (O, Ne, Fe, Mg, Si, and S; also N for AB Aur), while we fixed the abundances of those elements that do not show significant features (C, N, Ar, Ca, Al, and Ni; S for AB Aur) at the values used in the XEST data analysis (C=0.45, N=0.788, Ar=0.55, Ca=0.195, Al=0.5, Ni=0.195, and S=0.417; see Güdel et al. 2007a). All abundances are calculated with respect to the solar photospheric abundances of Anders & Grevesse (1989) except Fe, for which we used the value given by Grevesse & Sauval (1999). The above coronal abundances are characteristic for active, young stars described in the literature (García-Alvarez et al. 2005; Argiroffi et al. 2004). They are arranged according to a weak “inverse First Ionization Potential (FIP) effect” (elements with higher FIP are overabundant relative to low-FIP elements if normalized to solar photospheric abundances; Brinkman et al. 2001; Güdel et al. 2001a).

Finally, the X-ray luminosity $L_X$ was determined in the energy range between 0.3 and 10 keV from the best-fit model, assuming a distance of 140 pc (e.g., Loinard et al. 2005, Kenyon et al. 1994).

### 3.4 Results

#### 3.4.1 Light curves

In Fig. 3.1, the light curves of the nine stars are shown. The star V410 Tau was observed twice consecutively (XEST-23 and XEST-24). Both light curves are shown in our figure. For each light curve we also show the hardness, defined...
Figure 3.1: Light curves of the nine stars. Counts from all EPIC detectors have been combined, except for SU Aur and AB Aur, where data only from the two MOS detectors were available. Binning is to 1 ks, except for AB Aur, where 5 ks bins are used. Count rates refer to the energy range of 0.5-7.3 keV (except for AB Aur, where we used the range 0.3-4 keV). The size of the crosses in the x direction represent the bin width. Hardness is given by the ratio between the hard band (1-4 keV for AB Aur, 2-7.3 keV for all other stars) and the soft band (0.3-1 keV for AB Aur, 0.3-2 keV for all other stars). The dashed line in the light curve of V410 Tau separates the observations of XEST-23 and XEST-24. The dashed lines in the BP Tau light curve mark the flaring time interval excluded from the spectral fit. The dotted curve overplotted on the AB Aur light curve shows a sinusoidal fit.
3.4. Results

as the ratio between the hard band counts (1-4 keV for AB Aur, 2-7.3 keV for all other stars) and the soft band counts (0.3-1 keV for AB Aur, 0.3-2 keV for all other stars).

No strong variability (exceeding a factor of two between minimum and maximum in count rate) is seen in the light curves of V773 Tau, HP Tau/G2, and DN Tau. HD 283572 displays slow variability on timescales of 20-30 ks. The count rate varies by about 30%. Because the observation lasts about one third of the stellar rotation period, the variation could be due to rotational modulation of X-ray emitting regions in the corona, but equally well due to evolution of active regions.

DH Tau’s brightness decreases during the entire observation. Possibly, a large flare occurred before the observation. Further evidence for this hypothesis is the continuously decreasing hardness. Additional support for a flare during the XMM-Newton observation will be discussed in Sect. 3.4.5.

The light curve of the Herbig star AB Aur also shows variability. We have fitted the curve with a sine function and we find a modulation with a period of 42.2 hr. A similar modulation was also found in the Mg II lines that are thought to be formed in the wind of AB Aur (see Chapter 5 for a detailed discussion).

The two observations of V410 Tau were taken consecutively. The light curve from XEST-23 displays a flare, which has largely decayed in XEST-24. A small flare is visible at about 35 ks in the second light curve. For the spectral fit, we used only the second observation (XEST-24). For a more detailed study of the X-ray variability of this and other TMC X-ray sources, see Stelzer et al. (2007) and Franciosini et al. (2007). Finally, flares are observed in the light curves of SU Aur and BP Tau. SU Aur exhibits three flares superposed on a slowly decaying light curve. For BP Tau, the total contribution of the flare to the recorded counts is modest, and the quiescent emission contained a sufficient number of counts so that we excluded the flare for the spectral fit. The time interval of the observation ignored for the analysis is marked by the dotted lines (between 15 and 40 ks after the observation start).

3.4.2 Spectra

The EPIC MOS spectra of the nine target stars are shown in Fig. 3.2. For display purposes, we have multiplied the spectra with different factors (see figure caption). The effect of absorption is clearly visible on the low-energy slopes of the spectra. The latter are steep for SU Aur and HP Tau/G2, demonstrating high photoelectric absorption, while they are shallow in the other sources. The spectra of HD 283572, SU Aur, and marginally V773 Tau show the Fe line complex at 6.7 keV, which is a signature of very hot coronal plasma. We also note the presence of Ly α and He-like lines of Mg, Si, and S for these stars. In contrast, the spectrum of AB Aur falls off rapidly above 1 keV, suggesting a
Figure 3.2: EPIC MOS1 spectra of the nine target stars. For plotting purposes the spectra have been multiplied with different factors: $10^{-5}$ for DN Tau, $10^{-4}$ for BP Tau and DH Tau, $10^{-3}$ for HP Tau/G2, $10^{-2}$ for SU Aur, $10^{-1}$ for V410 Tau, 1 for V773 Tau, and 10 for HD 283572.

Fig. 3.3 shows the fluxed RGS spectra of our sample. The spectra of HD 283572, SU Aur, and HP Tau/G2 reveal a strong continuum clearly pointing to hot plasma. The spectra of SU Aur and HD 283572 show unusually high flux ratios between the strongest Fe lines and the NeIX or NeX lines when compared with other spectra, suggesting a higher relative Fe abundance in these two stars.

A rather strong O\textsc{vii} triplet is seen in the spectra of BP Tau, DN Tau, and AB Aur. In the other CTTS spectra, the triplet is not present, probably dominance of cool plasma.
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Figure 3.3: Co-added and fluxed spectra from the RGS1 and RGS2 instruments. The spectra are background subtracted and are binned to a bin width of 0.066 Å for HP Tau/G2, 0.058 Å for DN Tau and DH Tau, 0.050 Å for BP Tau, 0.042 Å for AB Aur, and 0.035 Å for the other stars. In each spectrum we also plot the typical 1σ errors at 14.5 Å and at 21.6 Å (at the position of the O vii line). The arrow in the HD 283572 spectrum indicates the precise wavelength of the O vii He/β line at 18.63 Å.
Figure 3.4: Data and best fit spectrum (EMD model) for V773 Tau. The best-fit model is shown by the histograms in the wavelength regions used for the fit.

because the spectra are more absorbed. Despite the low $N_H$ for WTTS, their spectra do not reveal the O\textsuperscript{vii} lines.\textsuperscript{1}

### 3.4.3 Thermal Structure

We now present results of our spectral fits. The fitted parameters are listed in Table 3.4 and Table 3.5 for the EMD model and the 2-$T$ or 3-$T$ model, respectively. In Fig. 3.4 we plot, as an example, the data and the best fit of the EMD model for V773 Tau; the fits to the other spectra are similar. The reduced $\chi^2$ are good for all fits ($\chi^2_{\text{red}} \lesssim 1.3$) and are very similar for the two approaches.

For each star and each model we computed the average temperature ($T_{\text{av}}$) as the logarithmic average of all temperatures used in the fit, applying the emission measures as weights. We see a wide range of thermal properties, from the very hot SU Aur with an average temperature of 20–23 MK, an EMD peaking at about 7.7 MK and a nearly flat high-$T$ slope ($\beta = -0.05$), to the cool AB Aur, with $T_{\text{av}} \approx 4.7$ MK and an EMD peaking at 4.4 MK with a steep high-$T$ slope ($\beta = -1.9$). For HP Tau/G2, DN Tau, and AB Aur, we

\textsuperscript{1}The RGS1 was not in use for the observation of HD 283572 and the RGS2 does not cover the O\textsuperscript{vii} triplet due to a CCD chip failure.
Table 3.4: Results from the spectral fits using the EMD model.\(^1\)

<table>
<thead>
<tr>
<th>Parameters</th>
<th>HD 283572</th>
<th>V 773 Tau</th>
<th>V 410 Tau</th>
</tr>
</thead>
<tbody>
<tr>
<td>(N_H) [10(^{22}) cm(^{-2})]</td>
<td>= 0.08 (^2)</td>
<td>= 0.17 (^2)</td>
<td>= 0.02 (^2)</td>
</tr>
<tr>
<td>(T_e) [MK]</td>
<td>10.20 (9.20, 11.13)</td>
<td>8.72 (7.71, 9.74)</td>
<td>13.66 (10.79, 15.17)</td>
</tr>
<tr>
<td>(EM) [10(^{52}) cm(^{-3})]</td>
<td>127.98</td>
<td>75.18</td>
<td>45.22</td>
</tr>
<tr>
<td>(\beta)</td>
<td>-0.84 (-0.99, -0.70)</td>
<td>-0.53 (-0.67, -0.40)</td>
<td>-1.59 (-1.80, -0.87)</td>
</tr>
<tr>
<td>O (^4)</td>
<td>0.24 (0.19, 0.29)</td>
<td>0.48 (0.39, 0.59)</td>
<td>0.39 (0.30, 0.47)</td>
</tr>
<tr>
<td>Ne (^4)</td>
<td>0.50 (0.40, 0.62)</td>
<td>1.30 (1.09, 1.55)</td>
<td>1.08 (0.83, 1.31)</td>
</tr>
<tr>
<td>Mg (^4)</td>
<td>0.51 (0.41, 0.63)</td>
<td>0.50 (0.37, 0.65)</td>
<td>0.26 (0.15, 0.38)</td>
</tr>
<tr>
<td>Si (^4)</td>
<td>0.25 (0.19, 0.32)</td>
<td>0.29 (0.21, 0.39)</td>
<td>0.12 (0.04, 0.20)</td>
</tr>
<tr>
<td>S (^4)</td>
<td>0.18 (0.08, 0.28)</td>
<td>0.64 (0.46, 0.83)</td>
<td>0.48 (0.33, 0.65)</td>
</tr>
<tr>
<td>Fe (^4)</td>
<td>0.33 (0.27, 0.39)</td>
<td>0.29 (0.23, 0.36)</td>
<td>0.20 (0.16, 0.24)</td>
</tr>
<tr>
<td>(T_{av}) [MK]</td>
<td>14.91</td>
<td>15.60</td>
<td>14.74</td>
</tr>
<tr>
<td>(L_X) [10(^{30}) erg/s](^5)</td>
<td>13.26</td>
<td>8.85</td>
<td>4.57</td>
</tr>
<tr>
<td>(\chi^2)</td>
<td>0.94</td>
<td>0.94</td>
<td>1.24</td>
</tr>
<tr>
<td>dof</td>
<td>251</td>
<td>208</td>
<td>173</td>
</tr>
<tr>
<td>Parameters</td>
<td>HP Tau/G2</td>
<td>SU Aur</td>
<td>DH Tau</td>
</tr>
<tr>
<td>(N_H) [10(^{22}) cm(^{-2})]</td>
<td>= 0.41 (^2)</td>
<td>= 0.32 (^2)</td>
<td>= 0.20 (^2)</td>
</tr>
<tr>
<td>(T_e) [MK]</td>
<td>10.06 (9.07, 11.41)</td>
<td>7.66 (6.95, 8.13)</td>
<td>14.77 (12.37, 17.12)</td>
</tr>
<tr>
<td>(EM) [10(^{52}) cm(^{-3})]</td>
<td>90.06</td>
<td>57.84</td>
<td>87.88</td>
</tr>
<tr>
<td>(\beta)</td>
<td>-1.15 (-1.39, -0.99)</td>
<td>-0.05 (-0.11, 0.06)</td>
<td>-2.34 (-3.00, -1.73)</td>
</tr>
<tr>
<td>O (^4)</td>
<td>0.20 (0.06, 0.38)</td>
<td>0.30 (0.24, 0.38)</td>
<td>0.49 (0.38, 0.64)</td>
</tr>
<tr>
<td>Ne (^4)</td>
<td>0.65 (0.43, 0.91)</td>
<td>0.38 (0.28, 0.53)</td>
<td>0.72 (0.55, 0.94)</td>
</tr>
<tr>
<td>Mg (^4)</td>
<td>0.51 (0.35, 0.70)</td>
<td>1.17 (1.05, 1.43)</td>
<td>0.33 (0.20, 0.48)</td>
</tr>
<tr>
<td>Si (^4)</td>
<td>0.21 (0.13, 0.31)</td>
<td>0.64 (0.57, 0.79)</td>
<td>0.16 (0.07, 0.25)</td>
</tr>
<tr>
<td>S (^4)</td>
<td>0.26 (0.12, 0.41)</td>
<td>0.57 (0.45, 0.72)</td>
<td>0.45 (0.30, 0.62)</td>
</tr>
<tr>
<td>Fe (^4)</td>
<td>0.43 (0.34, 0.56)</td>
<td>0.67 (0.61, 0.77)</td>
<td>0.16 (0.11, 0.23)</td>
</tr>
<tr>
<td>(T_{av}) [MK]</td>
<td>12.89</td>
<td>20.07</td>
<td>13.76</td>
</tr>
<tr>
<td>(L_X) [10(^{30}) erg/s](^5)</td>
<td>9.60</td>
<td>7.19</td>
<td>8.20</td>
</tr>
<tr>
<td>(\chi^2)</td>
<td>0.92</td>
<td>1.25</td>
<td>1.00</td>
</tr>
<tr>
<td>dof</td>
<td>131</td>
<td>446</td>
<td>142</td>
</tr>
<tr>
<td>Parameters</td>
<td>BP Tau</td>
<td>DN Tau</td>
<td>AB Aur</td>
</tr>
<tr>
<td>(N_H) [10(^{22}) cm(^{-2})]</td>
<td>= 0.06 (^2)</td>
<td>= 0.07 (^2)</td>
<td>= 0.05 (^2)</td>
</tr>
<tr>
<td>(T_e) [MK]</td>
<td>11.58 (10.03, 15.11)</td>
<td>6.31 (4.16, 9.08)</td>
<td>4.38 (2.69, 5.68)</td>
</tr>
<tr>
<td>(EM) [10(^{52}) cm(^{-3})]</td>
<td>11.24</td>
<td>15.37</td>
<td>5.12</td>
</tr>
<tr>
<td>(\beta)</td>
<td>-1.21 (-2.06, -0.93)</td>
<td>-0.66 (-1.06, -0.33)</td>
<td>-1.9 (-2.57, -1.52)</td>
</tr>
<tr>
<td>N (^4)</td>
<td>0.57 (0.30, 1.13)</td>
<td>0.57 (0.30, 1.13)</td>
<td>0.57 (0.30, 1.13)</td>
</tr>
<tr>
<td>O (^4)</td>
<td>0.46 (0.36, 0.59)</td>
<td>0.11 (0.05, 0.23)</td>
<td>0.22 (0.13, 0.32)</td>
</tr>
<tr>
<td>Ne (^4)</td>
<td>0.91 (0.66, 1.18)</td>
<td>0.45 (0.19, 0.84)</td>
<td>0.62 (0.48, 1.04)</td>
</tr>
<tr>
<td>Mg (^4)</td>
<td>0.50 (0.32, 0.69)</td>
<td>0.52 (0.27, 0.90)</td>
<td>0.28 (0.13, 0.74)</td>
</tr>
<tr>
<td>Si (^4)</td>
<td>0.26 (0.15, 0.38)</td>
<td>0.29 (0.12, 0.56)</td>
<td>0.90 (0.60, 1.32)</td>
</tr>
<tr>
<td>S (^4)</td>
<td>0.56 (0.34, 0.78)</td>
<td>0.11 (0.00, 0.56)</td>
<td>0.11 (0.00, 0.56)</td>
</tr>
<tr>
<td>Fe (^4)</td>
<td>0.18 (0.12, 0.24)</td>
<td>0.12 (0.06, 0.23)</td>
<td>0.29 (0.22, 0.47)</td>
</tr>
<tr>
<td>(T_{av}) [MK]</td>
<td>14.28</td>
<td>11.30</td>
<td>8.09</td>
</tr>
<tr>
<td>(L_X) [10(^{30}) erg/s](^5)</td>
<td>1.16</td>
<td>1.24</td>
<td>0.39</td>
</tr>
<tr>
<td>(\chi^2)</td>
<td>1.32</td>
<td>1.06</td>
<td>1.02</td>
</tr>
<tr>
<td>dof</td>
<td>161</td>
<td>45</td>
<td>79</td>
</tr>
</tbody>
</table>

---

1 68% error ranges are given in parentheses.
2 Held fixed at values found in the XEST survey (Güdel et al. 2007a).
3 Total EM integrated over temperature bins between \(\log T[K]= 6 – 7.9\) (see Güdel et al. 2007a for more details).
4 Element abundances are with respect to solar values given by (Anders & Grevesse 1989; Grevesse & Sauval 1999 for Fe).
5 Determined in the 0.3-10.0 keV band.
Table 3.5: Results from the spectral fits using the 2T/3T model.\(^1\)

<table>
<thead>
<tr>
<th>Parameters</th>
<th>HD 283572</th>
<th>V 773 Tau</th>
<th>V 410 Tau</th>
</tr>
</thead>
<tbody>
<tr>
<td>(N_\text{H} [10^{24} \text{ cm}^{-2}])</td>
<td>2.19 (1.33, 3.09)</td>
<td>4.51 (4.01, 5.77)</td>
<td>6.40 (3.96, 11.28)</td>
</tr>
<tr>
<td>(T_1 [\text{MK}])</td>
<td>8.60 (8.33, 9.08)</td>
<td>9.15 (8.42, 10.93)</td>
<td>9.77 (5.64, 14.17)</td>
</tr>
<tr>
<td>(T_2 [\text{MK}])</td>
<td>26.03 (24.96, 27.12)</td>
<td>29.39 (27.49, 32.46)</td>
<td>24.78 (22.71, 26.69)</td>
</tr>
<tr>
<td>(T_3 [\text{MK}])</td>
<td>18.77 (4.50, 47.13)</td>
<td>8.01 (4.29, 16.85)</td>
<td>8.61 (4.31, 17.16)</td>
</tr>
<tr>
<td>(EM_1 [10^{52} \text{ cm}^{-3}])</td>
<td>30.33 (26.00, 35.88)</td>
<td>20.28 (15.56, 22.43)</td>
<td>11.25 (6.06, 18.78)</td>
</tr>
<tr>
<td>(EM_2 [10^{52} \text{ cm}^{-3}])</td>
<td>65.79 (62.50, 69.00)</td>
<td>37.88 (34.30, 40.89)</td>
<td>23.32 (20.61, 26.61)</td>
</tr>
<tr>
<td>Parameters</td>
<td>BP Tau/672</td>
<td>SU Aur</td>
<td>DN Tau</td>
</tr>
<tr>
<td>(N_\text{H} [10^{24} \text{ cm}^{-2}])</td>
<td>1.20 0.99 1.06</td>
<td>0.48 (0.27, 0.68)</td>
<td>0.37 (0.27, 0.49)</td>
</tr>
<tr>
<td>(T_1 [\text{MK}])</td>
<td>8.86 (8.43, 9.19)</td>
<td>7.54 (7.30, 7.75)</td>
<td>13.97 (12.05, 18.06)</td>
</tr>
<tr>
<td>(T_2 [\text{MK}])</td>
<td>25.72 (23.42, 28.37)</td>
<td>38.14 (35.68, 41.14)</td>
<td>24.55 (21.36, 27.69)</td>
</tr>
<tr>
<td>(T_3 [\text{MK}])</td>
<td>35.41 (31.62, 40.19)</td>
<td>7.25 (4.53, 14.35)</td>
<td>37.93 (24.16, 45.46)</td>
</tr>
<tr>
<td>(EM_1 [10^{52} \text{ cm}^{-3}])</td>
<td>38.14 (35.68, 41.14)</td>
<td>24.55 (21.36, 27.69)</td>
<td></td>
</tr>
<tr>
<td>(EM_2 [10^{52} \text{ cm}^{-3}])</td>
<td>30.49 (24.66, 33.21)</td>
<td>34.79 (25.10, 46.39)</td>
<td></td>
</tr>
</tbody>
</table>

1 68% error ranges are given in parentheses.
2 Held fixed at values found in the XEST survey (Güdel et al. 2007a).
3 Element abundances are with respect to solar values given by (Anders & Grevesse 1989; Grevesse & Sauval 1999 for Fe).
4 Determined in the 0.3-10.0 keV band.
5 Determined in the 0.1-10.0 keV band.
fitted a 2-\(T\) model instead of the 3-\(T\) model because a third component was not required for a good fit.

Adding \(\alpha\) as a free fit parameter did not change the abundances significantly, but increased the error bars. The \(\chi^2\) values were nearly identical. The combination of fit parameters \(\alpha\), \(\beta\), and \(T_0\) was ill-constrained for our stellar spectra. This is a consequence of the weakness of our spectra, the small number of lines available, the breadth of the emissivity functions, and the interrelation between emission measure and abundances. We conclude that the EMD fit with fixed \(\alpha\) is satisfactory for our spectral analysis.

The average temperatures are similar for the two approaches. The results of the EMD fit are also in quite good agreement with the results from the EPIC fits (Table 3.3, after Güdel et al. 2007a). Furthermore, our results for SU Aur, BP Tau, and HD 283572 are also within the error ranges of the results reported by Robrade & Schmitt (2006) and Scelsi et al. (2005).

### 3.4.4 Abundances

The abundances are listed in Tables 3.4 and 3.5, and are shown graphically in Fig. 3.5 with respect to solar photospheric abundances. For AB Aur the coronal abundances normalized to the AB Aur photospheric values (Acke & Waelkens 2004) are also shown with the open circles. The abundance patterns are similar for the two adopted models. The abundances of HD 283572 first decrease for increasing FIP, reaching a minimum around 10 eV (element S). For higher FIP, the abundances increase with FIP. A similar trend, although less marked, can be observed for the abundances of HP Tau/G2.

The abundance patterns of V773 Tau, V410 Tau, DH Tau, and BP Tau are consistent with an inverse FIP effect. Similar patterns are observed in active stars and T Tauri stars (Chapter 2, Argiroffi et al. 2004). In contrast, the abundance pattern shown by SU Aur is peculiar among this sample: It is reminiscent of a solar-like FIP effect, i.e. elements with low FIP are more abundant than elements with larger FIP.

As suggested above, we find that the Fe abundances are larger in SU Aur, HP Tau/G2, and marginally in HD 283572 than in the rest of the sample. In the extreme case of SU Aur, the Fe abundance amounts to 0.67–0.81 times the solar photospheric value. Such high Fe abundance values have been reported for relatively inactive stars, while magnetically active stars usually show a strong Fe depletion (see Chapter 2 and Güdel 2004a).

The Ne/Fe abundance ratio reaches modest values for SU Aur (0.6-0.75), HD 283572 and HP Tau/G2 (1.5), and AB Aur (2.1-2.6). However, for the other T Tauri stars, regardless of their accretion state, the Ne abundance is 4-6 times higher than the Fe abundance. Such high Ne/Fe abundance ratios are unusual for main-sequence stars (in Chapter 2 we reported Ne/Fe below 2 for our sample of solar analogs), although they are reminiscent of ratios
Figure 3.5: Element abundances normalized to solar photospheric values (Anders & Grevesse 1989 and Grevesse & Sauval 1999 for Fe) derived from the EMD and 3T fits. Open circles for AB Aur: values normalized to the AB Aur photospheric abundances (Acke & Waelkens 2004).
3.4. Results

reported for RS CVn binaries (Ne/Fe = 5.3–13.4, excluding Capella for which Ne/Fe = 0.64, Audard et al. 2003a). A very high Ne/Fe abundance ratio has also been measured in TW Hya, reaching values of 7–11 (Kastner et al. 2002; J. Kastner, private communication; Stelzer & Schmitt 2004; B. Stelzer, private communication; referring to the solar photospheric abundances of Grevesse & Sauval 1999 for Fe), in TWA 5 (≈ 6, Argiroffi et al. 2005) and in HD 98800 (≈ 4, Kastner et al. 2004, J. Kastner, private communication). For BP Tau, a high Ne abundance has been reported before (Ne/Fe ≈ 5, Rorade & Schmitt 2006).

3.4.5 X-Ray Luminosities

Large changes in X-ray luminosity of coronal stars are usually accompanied by large changes in spectral parameters (e.g. temperature and densities), in particular in the case of flares. Therefore, it is important to understand if the observed $L_X$ is characteristic of the star as judged from previous observations. The X-ray luminosities from the two methods are in good agreement for all stars.

For HD 283572 $L_X \approx 1.3 \times 10^{31}$ erg s$^{-1}$ also agrees well with results found previously by Favata et al. (1998) using other X-ray telescopes ($L_X = [0.8 - 2.1] \times 10^{31}$ erg s$^{-1}$), and the value found by Scelsi et al. (2005, $L_X \approx 10^{31}$ erg s$^{-1}$).

For V773 Tau we measured $L_X \approx 8.8 \times 10^{30}$ erg s$^{-1}$. This value is slightly higher than that found by Feigelson et al. (1994) in their ROSAT observation ($L_X = 5.5 \times 10^{30}$ erg s$^{-1}$; in the energy range 0.2-2 keV). Our $L_X$ is consistent with the quiescent emission measured using ASCA by Skinner et al. (1997): they found $L_X = 1.23 \times 10^{31}$ erg s$^{-1}$ assuming a distance of 150 pc, corresponding to $L_X = 1.07 \times 10^{31}$ erg s$^{-1}$ at 140 pc.

Our $L_X$ for V410 Tau ($L_X = 4.6 \times 10^{30}$ erg s$^{-1}$ for XEST-24) is lower than the values found in the ROSAT observation by Strom & Strom (1994) ($L_X = 1.3\times10^{31}$ erg s$^{-1}$), while it agrees well with the highest value reported by Stelzer et al. (2003) for a set of recent Chandra observations ($L_X = [3.2 - 4.0] \times 10^{30}$ erg s$^{-1}$). This comparison suggests that the flare seen in the first part of our XMM-Newton observations has largely decayed in the second observation relevant for our study.

The $L_X$ of SU Aur ($L_X = [7.4 - 7.8] \times 10^{30}$ erg s$^{-1}$) is in agreement with the value measured by Robrade & Schmitt (2006) for the same observation ($L_X = 8.1 \times 10^{30}$ erg s$^{-1}$) and with the value found by Skinner & Walter (1998, $L_X = (8.4 \pm 0.09) \times 10^{30}$ erg s$^{-1}$) in the ASCA observation. For BP Tau, our $L_X$ ($\approx 1.2 \times 10^{30}$ erg s$^{-1}$) is lower than the value found by Robrade & Schmitt (2006, $L_X = 2.3 \times 10^{30}$ erg s$^{-1}$), probably because our $N_H$ is lower and we excluded the flare.
Overall, we find that the long-term variability illustrated by the above comparison is compatible with the short-term variations seen in our light curves. Variations within a factor of \( \approx 2 \) are common.

The DH Tau light curve is decreasing, suggesting that the source is decaying after a strong flare. To test the activity level of this source, we compare our XMM-Newton observation with ROSAT observations reported in the roshri catalog in W3Browse.\(^2\) DH Tau was detected twice by the ROSAT High-Resolution Imager (HRI) instrument, once with a count rate of 0.020 ct s\(^{-1}\) as 1RXH J042941.3+263256 in observation rh202636, and once as a fainter source in the wings of the nearby DI Tau (which was much fainter in the first observation) in observation rh201088 (combined source 1RXH J042942.6+263250). In this latter case, DH Tau was about four times fainter than DI Tau, and the total count rate of the two was 0.014 ct s\(^{-1}\). We used PIMMS\(^3\) to transform count rates to (unabsorbed) fluxes, adopting a temperature of 10 MK and \( N_H = 2 \times 10^{21} \) cm\(^{-2}\) (Güdel et al. 2007a) to find \( L_X = 2.3 \times 10^{30} \text{ erg s}^{-1} \) and \( L_X = 4 \times 10^{29} \text{ erg s}^{-1} \) in the two observations, i.e., 4-20 times lower than the \( L_X \) measured in our observation. We conclude that the light curve is the result of the decay of a large flare starting prior to our observation.

### 3.5 The He-like O\(^{\text{VII}}\) Triplet

The flux ratio of the forbidden (\( f \)) line at 22.1 Å and the intercombination (\( i \)) line at 21.8 Å of O\(^{\text{VII}}\) is density-sensitive in the electron-density range between \( 10^{10} \) cm\(^{-3}\) and \( 10^{12} \) cm\(^{-3}\) (Gabriel & Jordan 1969). In the case of high densities, the more frequent collisions trigger the excitation from the upper level of the forbidden transition, \( 1s2s \ \ ^3S_1 \), to the upper level of the intercombination transitions, \( 1s2p \ ^3P_1,2 \). As a consequence, the measured \( f/i \) ratio becomes smaller. Photoexcitation in an UV radiation field would also decrease the \( f/i \) ratio. The photon wavelength for the excitation would correspond to the energy difference of the two upper states, namely 1630 Å for the O\(^{\text{VII}}\) triplet. The UV radiation field is thus important for stars with \( T_{\text{eff}} \gtrsim 10^4 \) K, i.e. only for AB Aur in our sample.

The measured ratio \( R = f/i \) of the forbidden to intercombination line flux can be written as

\[
R = \frac{R_0}{1 + \phi/\phi_c + n_e/N_c} = \frac{f}{i} \tag{3.2}
\]

where \( R_0 \) is the limiting flux ratio at low densities and for O\(^{\text{VII}}\) has a value of \( R_0 \approx 3.85 \) (Blumenthal et al. 1972). \( N_c \) is the critical density at which

---

\(^2\)a service of the Exploration of the Universe Division at NASA/GSFC and the High Energy Astrophysics Division of the Smithsonian Astrophysical Observatory (SAO), http://heasarc.gsfc.nasa.gov/cgi-bin/W3Browse/w3browse.pl

\(^3\)http://heasarc.gsfc.nasa.gov/Tools/w3pimms.html
3.5. The He-like O\textvisiblespace{vii} Triplet

Figure 3.6: Fit of the O\textvisiblespace{vii} triplets using variable electron density. Left panel, BP Tau: The best fit for $n_e = 3.4 \times 10^{11}$ cm\textsuperscript{-3} is plotted as a solid histogram, while the lower density limit ($n_e = 6.4 \times 10^{10}$ cm\textsuperscript{-3}, 90 % confidence) is shown by the dotted line. Right panel, AB Aur: The solid histogram gives the best fit ($n_e \lesssim 10^{10}$ cm\textsuperscript{-3}), while the dotted histogram is for the 90% upper limit to $n_e$ ($n_e < 1.3 \times 10^{11}$ cm\textsuperscript{-3}).

$R = R_0/2$ ($N_e \approx 3.4 \times 10^{10}$ cm\textsuperscript{-3} for O\textvisiblespace{vii}, Blumenthal et al. 1972). The ratio $\phi/\phi_e$ is the radiation term needed for AB Aur.

In a thermal plasma, the flux of the resonance line, $r$, is larger than the flux of $f$. Furthermore, for plasma with temperatures higher than 1.5 MK, the sum $f + i$ is smaller than $r$, so that the “G ratio” $G = (f + i)/r$ is smaller than unity (Porquet et al. 2001). Given the low S/N ratio of our data in the wavelength range of interest, it is difficult to fit the triplet lines individually and simultaneously fulfill the constraints for the $R$ and $G$ ratios. We instead proceeded as follows: we used the best-fit results of the 3-$T$ model and then made use of the density-dependent calculations for the O\textvisiblespace{vii} line fluxes as implemented in the XSPEC vmekal code. We kept all parameters fixed, except for the electron density and the emission measure of the cooler component (to allow for slight adjustments of the total O\textvisiblespace{vii} line flux). Thus, the thermal structure intrinsic to the model sets the correct requirement for $G$, and simultaneously fixes the continuum. The electron density $n_e$ was then varied until a best fit for the fluxes of the O\textvisiblespace{vii} $i$ and $f$ lines was found. Only the wavelength region of interest was used for the fit (between 21.4 and 22.3 Å). The $R$ ratio was finally measured from the line fluxes.

In Fig. 3.6 we present the O\textvisiblespace{vii} triplet for the stars BP Tau, and AB Aur, the only two stars for which the triplet is clearly visible, together with the best fits and upper or lower limits to the densities. In the spectrum of DN Tau (Fig. 3.3) an excess of counts is present at the wavelengths of the O\textvisiblespace{vii}
triplet, but the S/N ratio is too low to fit the triplet.

For BP Tau, the background was particularly high near the O\textsc{vii} triplet. We therefore decided to further restrict the inclusion fraction of the cross-dispersion PSF to 70% \((xpsfincl = 70\%)\) specifically for this wavelength region only. The resulting spectrum thus contains fewer background counts and the triplet appears more clearly, obviously at the cost of some source counts. We also fitted the emission measure of the hottest component. This allows us to slightly adjust the continuum, that, because of slight background subtraction inaccuracies, was not accurately represented.

We used bins of 56 m\(\AA\) width for the fit. The best fit is represented by a solid histogram in the left panel of Fig. 3.6. We find a best-fit electron density of \(n_e = 3.4 \times 10^{11} \text{ cm}^{-3}\), corresponding to \(R = 0.23\). The dotted line represents the 90% lower limit, corresponding to \(n_{e,\text{min},90} = 6.4 \times 10^{10} \text{ cm}^{-3}\), and \(R = 1.07\). For the 68% error we find \(n_{e,\text{min},68} = 1.6 \times 10^{11} \text{ cm}^{-3}\), corresponding to \(R = 0.76\). Given the low flux in the O\textsc{vii} \(f\) line, we were unable to constrain upper limits to the densities. The best-fit density is in agreement with the densities found by Schmitt et al. (2005) and Robrade & Schmitt (2006), \(n_e = 3 \times 10^{11} \text{ cm}^{-3}\) and \(n_e = 3.2^{+3.5}_{-1.2} \times 10^{11} \text{ cm}^{-3}\), respectively.

In the right panel of Fig. 3.6, the O\textsc{vii} triplet of the Herbig star AB Aur is shown. Results for the electron density measured in AB Aur have been discussed in detail in Chapter 5. Here, we will report the main results in order to compare them with results from CTTS. We performed the fit again using a bin width of 56 m\(\AA\) to find an electron density below the low-density limit, \(n_e \lesssim 10^{10} \text{ cm}^{-3}\) and \(R = R_0\). The dotted line in Fig. 3.6b corresponds to the 90% upper limit, which is \(n_{e,\text{max},90} \approx (1.3 \pm 0.4) \times 10^{11} \text{ cm}^{-3}\) and \(f/i = 0.95\). For the 68% upper limit we found \(n_{e,\text{max},68} \approx (4.2 \pm 1.2) \times 10^{10} \text{ cm}^{-3}\), corresponding to \(f/i = 2.42\).

### 3.6 Discussion

#### 3.6.1 Abundance Patterns

The three stars with spectral type G (HD 283572, SU Aur, HP Tau/G2) have similar properties. They were classified by Herbig & Bell (1988) as SU Aurigae-like stars, i.e., late type F to K stars showing an H\(\alpha\) equivalent width smaller than 10 \(\AA\), weak Ca\textsc{ii} emission, very broad absorption lines (\(v\sin i > 50 \text{ km s}^{-1}\)), and a relatively high optical luminosity.

In X-rays, these stars display an abundance pattern with Fe enhanced relative to oxygen (Fe/O \(\approx 1.4\) to 2.2), in contrast to the later-type stars in our sample that show an inverse FIP effect (Fe/O \(\approx 0.35\) to 1.0). Furthermore, the Ne/Fe abundance ratio is relatively small for these three stars (0.5–1.5) if compared with the other T Tauri stars in our sample (Ne/Fe \(\approx 4–6\)). In
3.6. Discussion

Figure 3.7: Fe/O abundance ratio as a function of the Fe/Ne ratio for abundances derived using an EMD model (a) and a 2/3-T model (b). The K-M type stars are plotted in black, while the G-type stars are in red. The ZAMS 47 Cas B and EK Dra are plotted in blue for comparison in both plots (from Chapter 2). The T Tau star TWA 5, the post T Tau star PZ Tel, the ZAMS AB Dor, TW Hya1 and 2 are plotted in (a) for comparison (the analysis were based on an EMD method; see Argiroff et al. 2005, 2004; García-Alvarez et al. 2005; Kastner et al. 2002; Stelzer & Schmitt 2004). CR Cha and TW Hya3 have been added to (b) according to 3T fits reported by Robrade & Schmitt (2006).

In Fig. 3.7 a, b we plot the Fe/O abundance ratio as a function of the Fe/Ne ratio for the EMD and 2T/3T models, respectively. The two stellar populations are different: the K-M type stars (plotted in black) are lower in both Fe/O and Fe/Ne than the G-type stars (plotted in red). One possible explanation is that the abundances might be related to mass or to surface gravity. However, then we would expect V410 Tau and V773 Tau to show abundance ratios similar to the G-type stars, as their masses are close to the masses of the SU Aurigae-like stars in our sample, but this is not observed. Scelsi et al. (2005) have also found similar abundances (when scaled to the Fe abundances) for three different G-type stars with different masses at different evolutionary stages (HD 283572, EK Dra, and 31 Com). They concluded that surface gravity is not a determining factor for coronal abundance ratios.

In Fig. 3.7 we also show for comparison the abundance ratios of the TTS CR Cha (Robrade & Schmitt 2006), TW Hya (labeled “TW Hya1” from Kastner et al. 2002, “TW Hya2” from Stelzer & Schmitt 2004, and “TW Hya3” from Robrade & Schmitt 2006), and TWA 5 (Argiroff et al. 2005). The abundances from these previous works have been converted to the photospheric abundance normalization used here, i.e. Grevesse & Sauval (1999) for Fe and Anders & Grevesse (1989) for O and Ne. CR Cha is of spectral type K2. TW Hya is
K8, and TWA 5 is M1.5. CR Cha and TW Hya3 were interpreted with a 3T model and are therefore plotted in Fig. 3.7b, while TW Hya1, TW Hya2, and TWA 5 were interpreted with an EMD method and are therefore plotted in Fig. 3.7a. TWA 5 shows abundance ratios well compatible with our other K-type TTS, while CR Cha’s ratios are somewhat higher. We note, however, that CR Cha is an early K star. TW Hya shows an Fe/Ne ratio similar to other K stars (especially in the measurements from Kastner et al. 2002 and Robrade & Schmitt 2006) while its Fe/O ratio is high.

An abundance pattern similar to that of SU Aurigae-like stars was also observed in zero-age-main-sequence (ZAMS) stars with spectral type G (47 Cas B, EK Dra, see Chapter 2). For comparison, we therefore plot in Fig. 3.7a, b the abundance ratios that were derived from a detailed EMD reconstruction and a multi-thermal fit, respectively, for these two G-type ZAMS stars. All G-type stars, independently of the evolutionary stage, display similar abundance ratios.

Near zero-age main sequence K stars like PZ Tel (Argiroffi et al. 2004) and AB Dor (Güdel et al. 2001a; García-Alvarez et al. 2005) show a classical inverse FIP pattern similar to what we observe in our K-type T Tau stars. The abundance ratios for PZ Tel (Argiroffi et al. 2004) and AB Dor (García-Alvarez et al. 2005) are also shown in Fig. 3.7a. Again, the abundance ratios are similar to the K-type TTS, although Fe/Ne is somewhat higher. We note, however, that both stars have somewhat earlier spectral types than is typical for our sample of T Tau stars: PZ Tel is classified as a K0 V star (Houk 1978) and AB Dor is of spectral type K0-2 V (Vilhu et al. 1987).

Further, we have checked the abundance ratios found in active G-type and K-type main-sequence stars in the previous literature using the compilation of Güdel (2004a) and a few recent references as given in Table 3.6. For these stars as well, both the Fe/Ne and Fe/O abundance ratios are higher we summarize the abundance ratios for G and K-type stars. The last three entries in Table 3.6 are stars that cannot be classified easily. AR Lac is composed of a G and K star and both components contribute strongly to the X-ray emission (Siarkowski et al. 1996). The source λ And is a G8 giant star, i.e. intermediate between the two samples, in fact showing abundances similar to K-type stars. Finally, AB Aur is a Herbig star. These last three stars are not included in the calculations of averages below. We do not report the errors because we are mainly interested in studying the distributions, i.e. their means and their standard deviations. Further, the abundances reported in the table originate from different works based on different methods, implying that error estimates may not be consistent with each other. Also, errors are not given in some papers. The mean ratios for G stars are \( \langle \text{Fe/Ne} \rangle_G = 1.02 \) (standard deviation \( \sigma = 0.48 \)) and \( \langle \text{Fe/O} \rangle_G = 2.03 \) (\( \sigma = 0.42 \)). For K stars (using the mean of the three abundance ratios for TW Hya) we find \( \langle \text{Fe/Ne} \rangle_K = 0.22 \) (\( \sigma = 0.11 \)) and \( \langle \text{Fe/O} \rangle_K = 0.58 \) (\( \sigma = 0.32 \)), i.e. substantially lower than for G-type stars.
### 3.6. Discussion

Table 3.6: Abundance ratios of active stars from this work and from the literature

<table>
<thead>
<tr>
<th>Star</th>
<th>Type</th>
<th>Spec.Type</th>
<th>Fe/Ne</th>
<th>Fe/O</th>
<th>Ne/O</th>
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<tbody>
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<td></td>
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<tr>
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<td>WTTS</td>
<td>G5</td>
<td>0.66</td>
<td>1.38</td>
<td>2.08</td>
<td>this work</td>
</tr>
<tr>
<td>HP Tau/G2</td>
<td>WTTS</td>
<td>G0</td>
<td>0.66</td>
<td>2.15</td>
<td>3.25</td>
<td>this work</td>
</tr>
<tr>
<td>SU Aur</td>
<td>CTTS</td>
<td>G2</td>
<td>1.76</td>
<td>2.23</td>
<td>1.27</td>
<td>this work</td>
</tr>
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<td>47 Cas</td>
<td>ZAMS</td>
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<td>1.57</td>
<td>2.74</td>
<td>Chapter 2</td>
</tr>
<tr>
<td>EK Dra</td>
<td>ZAMS</td>
<td>G0</td>
<td>1.26</td>
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<td>2.0</td>
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<td>Robrade &amp; Schmitt (2006)</td>
</tr>
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<td>K8</td>
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<td>1.44</td>
<td>10.06</td>
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<td>0.49</td>
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<td>RS CVn</td>
<td>K3-4</td>
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<td>0.45</td>
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<td>Other stars</td>
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<td>Herbig</td>
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</tr>
<tr>
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<td>RS CVn</td>
<td>K0+G2</td>
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<td>1.23</td>
<td>2.68</td>
<td>Huenemoerder et al. (2003)</td>
</tr>
<tr>
<td>λ And</td>
<td>RS CVn</td>
<td>G8</td>
<td>0.19</td>
<td>0.57</td>
<td>3.00</td>
<td>Audard et al. (2003a)</td>
</tr>
</tbody>
</table>

1 For TW Hya, the mean of the three abundance ratios (Fe/Ne resp. Fe/O) was used.
2 Excluding TW Hya.

Considering only the subsample of TTS, we find \( \langle \text{Fe/Ne} \rangle_{G,TTS} = 1.03 \) (standard deviation \( \sigma = 0.64 \)), \( \langle \text{Fe/O} \rangle_{G,TTS} = 2.92 \) (\( \sigma = 0.47 \)), \( \langle \text{Fe/Ne} \rangle_{K,TTS} = 0.23 \) (\( \sigma = 0.10 \)), and \( \langle \text{Fe/O} \rangle_{K,TTS} = 0.71 \) (\( \sigma = 0.35 \)), in agreement with the values found for the full sample. The Ne/O abundance ratios are also listed in Table 3.6. We find Ne/O to range between 1 and 3. Drake et al. (2005b) suggested that this ratio might be sensitive to accretion; however, we find no difference in the Ne/O ratio between CTTS (\( \langle \text{Ne/O} \rangle_C = 2.24 \), \( \sigma_C = 1.12 \),
excluding TW Hya) and WTTS ($\langle\text{Ne/O}\rangle_W = 2.75$, $\sigma_C = 0.43$). The mean Ne/O ratio for all (K and G-type) stars in Table 3.6 (excluding TW Hya) is $\langle\text{Ne/O}\rangle = 2.76$ ($\sigma = 1.65$). The only star with exceptionally high Ne/O ratio remains TW Hya. This might be related to the older age of TW Hya and the consequent evolution of grains in the disk (Drake et al. 2005b). Günther et al. (2006) recently reported an anomalously high Ne/O ratio also for the CTTS binary V4046 Sgr.

We conclude that a separation is visible between G-type stars and mid-K-M-type stars, with G stars having a larger Fe/Ne abundance ratio. Early K-type stars (like AB Dor, PZ Tel and CR Cha) show an intermediate Fe/Ne ratio. A separation also exists in the Fe/O abundance ratio if we exclude TW Hya. This star is, however, peculiar among TTS, since almost only cool plasma is present and the abundances refer, in contrast to other TTS, essentially to this cool plasma. It seems therefore that the abundance pattern in the coronae of pre-main-sequence and near-ZAMS stars relates to the spectral type, i.e. is a function of the photospheric temperature. We caution that our sample is small, and more studies are needed to consolidate this trend.

### 3.6.2 A Soft Excess in Accreting Stars

In the RGS spectra of Fig. 3.3, we notice that the weakly absorbed spectra of accreting stars (BP Tau, DN Tau, and AB Aur) display a relatively strong O\text{vii} triplet when compared with the O\text{viii} Ly$\alpha$ line. On the contrary, the O\text{vii} triplet is not visible in the spectrum of V410 Tau (also subject to weak absorption, $N_{\text{H}} = 2 \times 10^{20}$ cm$^{-2}$) This lets us hypothesize that a substantial cool plasma component is present in the accreting stars but not in the WTTS.

The ratio between the fluxes in the O\text{vii} triplet and the O\text{viii} line varies with temperature in the range of $\approx 1 - 5$ MK. In order to estimate this ratio, we derived the number of counts at the wavelengths where the lines are formed (between 18.75-19.2 Å for O\text{viii} and 21.4-22.2 Å for O\text{vii}). From the total number of counts measured in these wavelength intervals, we subtracted the number of background counts (scaled to the source area) and the number of counts due to continuum (derived from the EMD best-fit model). In the RGS1 the effective area drops substantially at wavelengths slightly shorter than O\text{viii}, and we therefore used the RGS2 spectrum to derive the number of counts in this usually bright line (we used the RGS1 spectrum only for AB Aur, because in this spectrum the effective area of RGS2 drops at wavelengths close to the O\text{viii} line). To obtain the fluence in the lines we divided the source counts by the effective areas at the relevant wavelengths. The number of counts and the fluences derived with this method are summarized in Table 3.7. In the last column, we also list the probability that the measured number of counts is due to Poissonian fluctuations in the background+continuum. The RGS2 spectrum of DN Tau is difficult to interpret quantitatively because it shows
3.6. Discussion

Figure 3.8: Ratio of the fluxes measured in the O\textsc{vii} triplet and the O\textsc{viii} Ly\alpha line as a function of $N_H$. The sample of young stars presented in this chapter is plotted with circles, while values taken from the literature are plotted with squares. The diamonds represent the main-sequence solar analogs of Chapter 2. HD 283572 is plotted with a triangle because the O\textsc{vii} line counts are not measured but derived from the best fit model. Filled symbols denote accreting stars, open symbols are WTTS. Errors are 1 $\sigma$.

some excess in the background around the O\textsc{viii} line, resulting in too low flux in the line. We therefore use the RGS1 to derive the the counts in the O\textsc{viii} line.

The O\textsc{vii}/O\textsc{viii} fluence (or equivalently, flux) ratio is plotted in Fig. 3.8 as a function of $N_H$. The dotted lines show the theoretical O\textsc{vii}/O\textsc{viii} ratio for isothermal plasma at a given temperature as a function of $N_H$. The flux ratios measured in our sample are displayed by circles, while square symbols refer to measurements taken from the literature (Robrade & Schmitt 2006 for TW Hya, BP Tau, and CR Cha; Argiroffi et al. 2005 for TWA 5). For HD 283572, because the RGS1 was not available and the RGS2 does not cover
the O\textsuperscript{vii} wavelength region, we derived the flux ratio from the best-fit model. Cool temperatures can be determined, apart from the O\textsuperscript{vii} He\alpha line, by the O\textsuperscript{vii} He\beta line at 18.63 Å, but the line has not significantly been detected. The feature that appears in the HD 283572 spectrum in Fig. 3.3 is located slightly but significantly longward of the O\textsuperscript{vii} He\beta line (at 18.75 Å instead of 18.63 Å) while the O\textsuperscript{viii} Ly\alpha line is located at its laboratory wavelength. The excess in flux at 18.75 Å in this spectrum is due essentially to a single bin at 18.75 ± 0.03 Å, 2σ above the continuum, while the spectral fit represents the data at the correct line wavelength accurately, implying a low flux also for the O\textsuperscript{vii} He\alpha lines. For a further check, we studied a spectrum of HD 283572 observed by Chandra (Audard et al. 2007, in preparation). This spectrum shows a well-developed O\textsuperscript{viii} line but a line neither at 18.62 Å nor at 18.75 Å. The 95% confidence upper limit for the presence of excess flux at 18.62 Å is approximately 10% of the O\textsuperscript{viii} Ly\alpha flux. Because the emissivity of the O\textsuperscript{vii} He\beta line is about 14% of the emissivity of the O\textsuperscript{vii} He\alpha r line at 21.6 Å and the latter is about 60% of the total triplet flux under typical conditions, we conclude that a flux in the O\textsuperscript{vii} triplet must be lower than O\textsuperscript{viii} Ly\alpha line flux at the 95% confidence level. In any case, a strong line at 18.6-18.8 Å as tentatively suggested by the XMM-Newton RGS spectrum can be excluded. The Chandra spectrum shows no indication of flux in the region of the O\textsuperscript{vii} triplet, but the effective area of the HETGS instrument used in this observation is too small to be useful for our study, because even a triplet with a total flux equal to the flux in the O\textsuperscript{viii} Ly\alpha line would not have been detected. For SU Aur and HP Tau/G2, the counts in the O\textsuperscript{vii} triplet were very close to zero, and therefore only 95% upper limits to the flux ratios are shown. We also note that these stars show the highest $N_{\text{H}}$ in our sample, which is the reason for strong suppression of the O\textsuperscript{vii} triplet. Although the triplet is not explicitly visible in the spectra of V410 Tau, V773 Tau, and DH Tau, we measured a slight excess of counts in the relevant wavelength interval. For these stars, we therefore plot their O\textsuperscript{vii}/O\textsuperscript{viii} flux ratios at their best-fit loci.
Table 3.7: Number of counts and fluences for the O\textsuperscript{viii} and O\textsuperscript{vii} lines. Cts(tot), Cts(bkg), Cts(cont), and Cts(src) are the number of counts measured in the total spectrum, in the background spectrum, in the continuum (computed using the EMD fit results), and in the line (Cts[tot]-Cts[bkg]-Cts[cont]). The last column give the probability that the measured number of counts is due to fluctuations in the background and continuum.

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<th>Star</th>
<th>Line</th>
<th>Spectrum used</th>
<th>Cts(tot)</th>
<th>Cts(bkg)</th>
<th>Cts(cont)</th>
<th>Cts(line)</th>
<th>Eff. Area [cm\textsuperscript{2}]</th>
<th>Fluence [ph/cm\textsuperscript{2}]</th>
<th>Prob.</th>
</tr>
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<td>O\textsuperscript{viii}</td>
<td>RGS2</td>
<td>81</td>
<td>10.57</td>
<td>11.53</td>
<td>58.90 ± 9.13</td>
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<td>O\textsuperscript{vii}</td>
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<td>RGS2</td>
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<td>1.38 ± 0.18</td>
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<td>RGS2</td>
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<td>O\textsuperscript{vii}</td>
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<td>RGS2</td>
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<td>6.66</td>
<td>4.32</td>
<td>9.02 ± 4.62</td>
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<td>RGS1</td>
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<td>6.02</td>
<td>38.62 ± 9.16</td>
<td>38</td>
<td>0.98 ± 0.25</td>
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<td>RGS2</td>
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<td>–</td>
<td>33.07</td>
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<td>52.5</td>
<td>1.60 ± 0.21</td>
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<tr>
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<td>O\textsuperscript{vii}</td>
<td>RGS1</td>
<td>≈ 55</td>
<td>–</td>
<td>36.51</td>
<td>18.18 ± 7.40</td>
<td>41</td>
<td>0.44 ± 0.18</td>
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</table>
For all WTTS, marked with open symbols in Fig. 3.8, we measure a low \( \text{O}^\text{vii}/\text{O}^\text{viii} \) ratio, even if \( N_\text{H} \) is small \( (2 \times 10^{20} \, \text{cm}^{-2} \) for V410 Tau). The temperatures corresponding to these line ratios \( (T_{\text{oxy}}) \) are consistent with an isothermal plasma of \( > 3.5 \, \text{MK} \). The lack of strong \( \text{O}^\text{vii} \) is evident in Fig. 3.3.

If the total flux in the \( \text{O}^\text{vii} \) were similar to the flux in the \( \text{O}^\text{viii} \) Ly\( \alpha \) line, then the \( \text{O}^\text{vii} \) \( r \)-line would be approximately 50-60\% of the \( \text{O}^\text{viii} \) Ly\( \alpha \) flux if \( N_\text{H} \) is low, which is not seen in Fig. 3.3 for these WTTS.

For comparison, we also plot the \( \text{O}^\text{vii}/\text{O}^\text{viii} \) ratios of the six solar analog stars presented in Chapter 2. These main sequence stars are almost unabsorbed; for illustration purposes, we plot them at \( N_\text{H} \geq 0.01 \times 10^{22} \, \text{cm}^{-2} \). The \( \text{O}^\text{vii}/\text{O}^\text{viii} \) flux ratios (or the upper limits thereof) that we measure in WTTS compare well with the same ratios measured in active ZAMS stars such as 47 Cas B or EK Dra. With regard to the cool end of the coronal emission measure distribution, WTTS and ZAMS stars seem to behave similarly.

On the other hand, we measure a high \( \text{O}^\text{vii}/\text{O}^\text{viii} \) only for accreting stars. For most of them, the line ratio is consistent with \( T_{\text{oxy}} \approx 2.5 - 3 \, \text{MK} \). These ratios are reminiscent of \( \text{O}^\text{vii}/\text{O}^\text{viii} \) found in rather inactive, evolved solar analogs. In contrast to the CTTS, however, the coronae of those more evolved stars are dominated by cool plasma, while much hotter plasma is common in CTTS.

For two CTTS, namely SU Aur and DH Tau, we measure an \( \text{O}^\text{vii}/\text{O}^\text{viii} \) ratio lower than for the other accreting stars. However, we have noticed that both stars are flaring (see Sect. 3.4.1), which must lower their \( \text{O}^\text{vii}/\text{O}^\text{viii} \) flux ratio. The \( \text{O}^\text{viii} \) line is sensitive to the hot plasma, so that its flux increases while the source is flaring. In contrast, the \( \text{O}^\text{vii} \) triplet is insensitive to the high temperatures measured in a strong flare. We therefore expect the \( \text{O}^\text{vii}/\text{O}^\text{viii} \) flux ratio to decrease if the source is flaring.

\( N_\text{H} \) does alter the \( \text{O}^\text{vii}/\text{O}^\text{viii} \) ratio, but \( N_\text{H} \) cannot be made responsible for the lack of \( \text{O}^\text{vii} \) flux detected in some stars. The range of \( N_\text{H} \) measured in the WTTS sample is in fact similar to the range measured in CTTS.

According to the \( \text{O}^\text{vii}/\text{O}^\text{viii} \) ratio measured in the sample investigated here, it is therefore possible that a soft excess is present in all accreting stars. Güdel et al. (2007b) presented the spectrum of the CTTS T Tau. Although the corona of this star is extremely hot, a soft plasma component must be present in order to explain the strong \( \text{O}^\text{vii} \) flux. The flux ratio measured in this star is \( \text{O}^\text{vii}/\text{O}^\text{viii} = 1.06 \pm 0.29 \) for \( N_\text{H} = 0.48 \times 10^{22} \, \text{cm}^{-2} \). The latter measurements confirm the presence of a soft excess in T Tau, consistent with the results that we find here. A specific discussion is presented by Güdel et al. (2007b).

In order to illustrate the above trend using physical properties that are not biased by other stellar properties, we derive the temperature \( T_{\text{oxy}} \) using the loci for isothermal plasma plotted in Fig. 3.8. \( T_{\text{oxy}} \) for each star was computed using a spline interpolation. The results are shown in Table 3.8. \( T_{\text{oxy}} \) of V410
3.6. Discussion

Table 3.8: $T_{\text{oxy}}$ derived from Fig. 3.8 for each star.

<table>
<thead>
<tr>
<th>Star</th>
<th>$T_{\text{oxy}}$</th>
</tr>
</thead>
<tbody>
<tr>
<td>BP Tau</td>
<td>2.97 (2.77 3.36)</td>
</tr>
<tr>
<td>V773 Tau</td>
<td>7.0 (4.14 7.0)</td>
</tr>
<tr>
<td>V410 Tau</td>
<td>7.0 (7.0 7.0)</td>
</tr>
<tr>
<td>HP Tau</td>
<td>$&gt; 2.36$</td>
</tr>
<tr>
<td>SU Aur</td>
<td>$&gt; 3.10$</td>
</tr>
<tr>
<td>DH Tau</td>
<td>3.95 (3.24 7.0)</td>
</tr>
<tr>
<td>DN Tau</td>
<td>2.84 (2.35 4.66)</td>
</tr>
<tr>
<td>AB Aur</td>
<td>2.75 (2.47 3.11)</td>
</tr>
<tr>
<td>HD 2835721</td>
<td>4.51 (3.88 7.0)</td>
</tr>
</tbody>
</table>

Figure 3.9: Temperature corresponding to the observed O\text{vii}/O\text{viii} count ratio as a function of H\alpha equivalent width (left) and as a function of the accretion rate $\dot{M}$ (right). Errors in log $T_{\text{oxy}}$ are 1σ.

Tau and V773 Tau are approximately at or below the loci for 7 MK, above which the O\text{vii}/O\text{viii} ratio is no longer sensitive to temperature. We therefore assign a lower limit for $T_{\text{oxy}}$ of 7 MK to both stars. In Fig. 3.9, we correlate the $T_{\text{oxy}}$ with stellar accretion parameters. Filled and open circles represent the CTTS and WTTS, respectively.

In the left panel of Fig. 3.9, we plot $T_{\text{oxy}}$ as a function of the H\alpha equivalent width. For the stars in the XEST survey, the values of EW(H\alpha) are taken from Gudel et al. (2007a) and references therein (see Table 3.1). The EW(H\alpha) values for TW Hya and TWA 5 are from Reid (2003), namely 220 Å and 13.4 Å, respectively, while the value for CR Cha is taken from Gauvin & Strom (1992), namely 43.6 Å. For HD 283572, an equivalent width of 0 Å has been reported (Kenyon et al. 1998); we plot an upper limit at EW (H\alpha) = 1.5 Å for...
illustration purposes. Generally, we find that stars with a large EW(Hα) show a low $T_{\text{oxy}}$. One problem with using EW(Hα) as an accretion signature is that for the same accretion rate ($\dot{M}$), EW(Hα) is smaller in G-type stars than in K-type stars due to formers’ high continuum.

We therefore plot in the right panel of Fig. 3.9 $T_{\text{oxy}}$ as a function of $\dot{M}$ (Table 3.1, Güdel et al. 2007a). In the cases of HD 283572 and HP Tau/G2, for which no upper limit for the accretion rate has been reported, we assign an upper limit of $\dot{M} < 10^{-9} \, M_\odot \, \text{yr}^{-1}$ because for accreting stars with masses similar to HD 283572 and HP Tau/G2, accretion rates between $10^{-9}$ and $10^{-8} \, M_\odot \, \text{yr}^{-1}$ were clearly identified (Calvet et al. 2004). For V773 Tau and V410 Tau we used upper limits for the accretion rate as reported in the XEST catalog (see Güdel et al. 2007a for a summary, White & Ghez 2001). TW Hya is plotted twice, as two different accretion rates have been reported (see Kastner et al. 2002): from Hα measurements, Muzerolle et al. (2000) derived $\dot{M} \approx 5 \times 10^{-10} \, M_\odot \, \text{yr}^{-1}$, while according to the excess in the CIV λ1549 line and the empirical relation between this excess and the accretion rate described by Johns-Krull et al. (2000), $\dot{M}$ could reach values of $(3 - 6) \times 10^{-8} \, M_\odot \, \text{yr}^{-1}$. We found no information on an upper limit of $\dot{M}$ for TWA 5 in the published literature, and again recall that this stellar system contains four components.

Despite the low statistics, a trend is clearly visible: stars with low (or vanishing) accretion rates have $T_{\text{oxy}}$ higher than the accreting stars. Observationally, these stars thus reveal a low Ovii/Oviii flux ratio.

### 3.6.3 X-Rays from Accretion?

Kastner et al. (2002) and Stelzer & Schmitt (2004) have interpreted the low $f/i$ flux ratio (consistent with high electron density), and other properties of the X-ray spectrum of TW Hya, such as the low X-ray temperature and the low Fe/Ne abundance ratio, in terms of accretion. An analogous scenario has been suggested by Schmitt et al. (2005), Robrade & Schmitt (2006), and Günther et al. (2006) to explain the low $f/i$ ratio measured in three other stars, BP Tau, CR Cha, and V4046 Sgr. In the last two cases, however, the accretion shock would only be responsible for the softest component of the spectrum, while magnetic activity or star-disk interaction is required to explain the hotter components.

In our stellar sample, an Ovii triplet can be measured for only two stars: the CTTS BP Tau and the Herbig star AB Aur. For the two accreting stars SU Aur and DH Tau, the Ovii triplet is not present, presumably because it is more absorbed ($N_H = [3.1 - 3.2] \times 10^{21} \, \text{cm}^{-2}$ and $N_H = 2.0 \times 10^{21} \, \text{cm}^{-2}$, respectively) and perhaps also because the X-ray emission was dominated by flaring plasma.

The low $f/i$ flux ratio measured in BP Tau is consistent with high densities ($n_e = 3.4 \times 10^{11} \, \text{cm}^{-3}$). On the other hand, the $f/i$ ratio measured in AB Aur
strictly excludes high densities \((n_e \lesssim 10^{10} \text{ cm}^{-3})\).

We now test the accretion hypothesis for BP Tau and AB Aur. The accretion rate of BP Tau is \(M_{\text{acc}} \approx (1.32 - 2.88) \times 10^{-8} \text{M}_\odot \text{ yr}^{-1}\) (White & Ghez 2001; Muzerolle et al. 1998). For AB Aur we approximate the accretion rate as \(\approx 10^{-8} \pm 1 \text{M}_\odot \text{ yr}^{-1}\), according to Chapter 5 and references therein. The accretion luminosity, assuming that the disk is truncated at the corotation radius \((R_{\text{cor}})\), is given by \(L_{\text{acc}} = (1 - R/R_{\text{cor}})GM\dot{M}/R\), which can be written as \(L_{\text{acc,30}} \approx 1200(1 - \hat{R}/R_{\text{cor}})M M_{-8}/R\), where \(L_{\text{acc,30}} = L_{\text{acc}}/(10^{30} \text{ erg s}^{-1})\), \(M = M/\text{M}_\odot\), \(M_{\text{acc,-8}} = M_{\text{acc}}/10^{-8} \text{M}_\odot \text{ yr}^{-1}\), \(\hat{R} = R/R_\odot\), and \(R_{\text{cor}} = R_{\text{cor}}/R_\odot\). Using the stellar parameters of Table 3.1 we obtain, \(R_{\text{rms cor, BP}} \approx 7.5 R_{\text{BP}}\) and \(R_{\text{cor, AB}} \approx 2 R_{\text{AB}}\), and X-ray luminosities \(L_{\text{acc, BP}} \gtrsim 3 \times 10^{32} \text{ erg s}^{-1}\) and \(L_{\text{acc, AB}} \approx 7 \times 10^{32} \text{ erg s}^{-1}\), i.e. enough to account for the observed luminosities.

We further estimate the temperature expected for the accretion shock. The temperature in case of strong shocks is given by \(T = 3v^2 \mu m_p/16k\), where the velocity is approximately the free fall velocity \(v_{\text{ff}} = (2GM/R)^{1/2}\), \(m_p\) is the proton mass, \(k\) is the Boltzmann constant, and the mean molecular weight \(\mu \approx 0.62\) for a fully ionized gas. We thus find \(T \approx 5.4 \times 10^{6} \dot{M}/R \text{ [K]}\), and with the parameters from Table 3.1, \(T_{\text{BP}} \approx 2 \text{ MK}\) and \(T_{\text{AB}} \approx 6 \text{ MK}\). For BP Tau, this temperature is consistent with the temperature of the soft component in the 3-T fit and with the temperature derived from Fig. 3.8; for AB Aur, \(T_{\text{AB}}\) is consistent with \(T_{\text{av}}\).

We can further estimate the shock density, using the strong-shock condition \(n_2 = 4 n_1\), where \(n_1\) and \(n_2\) are the pre-shock and post-shock densities. The density \(n_1\) can be estimated from the accretion mass rate and the accreting area on the stellar surface: \(\dot{M} \approx 4\pi R^2 f v_{\text{ff}} n_e m_p\), where \(f\) is the surface filling factor of the accretion flow. We thus find

\[
n_2 \approx \frac{4 \times 10^{11} M_{-8}}{R^{3/2} f} \text{ [cm}^{-3}]\].
\]

According to Calvet & Gullbring (1998), typical values for \(f\) are \(f = 0.1 - 0.10\). The density should then be \(n_2 = 10^{12} - 10^{14} \text{ cm}^{-3}\) for both stars, given the adopted \(\dot{M}\).

These densities are therefore compatible with the O vii triplet fluxes measured in BP Tau, but not with those in AB Aur. In order to obtain the electron density that we measure in AB Aur from Eq. 3.3, \(M_{\text{AB}}\) should be lowered to about \(10^{-10} \text{M}_\odot \text{ yr}^{-1}\); also the accreting area should be at least 10% and probably more, essentially the whole stellar surface. The first possibility is not supported from the (tentative) measurements of \(\dot{M}\) (Chapter 5 and references therein), while a filling factor approaching unity is unreasonable, given that the star accretes from a disk and a wind is present (Praderie et al. 1986). We also note that a low \(f/i\) ratio has also been measured for the CTTS T Tau (Güdel 2006a; Güdel et al. 2007b).
A problem with the accretion scenario is that the shock is formed close to the stellar photosphere (Calvet & Gullbring 1998) and that the X-rays could therefore be absorbed. Drake (2005a) studied the problem for TW Hya. The depth of the shock can be estimated from the measured electron density. For TW Hya the density of \( n_e \approx 10^{13} \text{ cm}^{-3} \), derived from Ne IX and O VII triplets, corresponds to a larger \( N_H \) than observed. Drake (2005a) suggested that \( n_e \) is only \( \approx 10^{12} \text{ cm}^{-3} \) in TW Hya while photoexcitation from the ambient UV radiation field could be responsible for the observed \( f/i \) ratio. For BP Tau, we measure an electron density smaller than in TW Hya, and we expect therefore that the shock is higher in the photosphere or above the photosphere, and the problem of photospheric absorption is at least in part alleviated, while the influence of the UV radiation field from the shock on the \( f/i \) ratio still remains unknown.

Accretion can therefore explain the soft excess that we measure in BP Tau, but cannot explain the soft X-ray emission in AB Aur.

### 3.7 Conclusions

We have presented high-resolution X-ray spectra of nine young stellar objects. Five of them are accreting stars (four CTTS and one Herbig star) and four are WTTS. From previous work on high-resolution X-ray spectroscopy of T Tauri stars (Kastner et al. 2002; Stelzer & Schmitt 2004; Argiroffi et al. 2005; Schmitt et al. 2005; Robrade & Schmitt 2006; Günther et al. 2006), three X-ray properties have been proposed to characterize accreting pre-main-sequence stars: i) they show strong soft emission (TW Hya); ii) they display a high electron density (TW Hya, BP Tau, V4046 Sgr); iii) the Ne or N abundances are relatively high (TW Hya, BP Tau, V4046 Sgr), when compared with Fe. However, the sample of stars that had been studied so far was too small to prove that these three properties are common to all CTTS. Our sample adds a significant number of spectra to test these conjectures.

For two accreting stars (BP Tau and AB Aur), we have been able to measure the O VII triplet and to derive the source electron density. While we measured a high density for BP Tau (3.4×10^{11} \text{ cm}^{-3}, confirming previous reports, Schmitt et al. 2005), the density for the Herbig star AB Aur is low, with \( n_e < 10^{10} \text{ cm}^{-3} \). In the high-resolution X-ray spectrum of the CTTS T Tau (XEST-01-045), the O VII triplet is also consistent with a low density, as reported by Güdel (2006a). The low signal-to-noise ratio does not allow to constrain the electron density for the CTTS CR Cha (Robrade & Schmitt 2006) and DN Tau sufficiently well. For the other CTTS, the O VII triplet was not measurable because of strong absorption or, probably, due to the dominance of hot material during strong flaring (in SU Aur and DH Tau).

None of the WTTS in our sample displays an O VII triplet, testifying to
the low flux in these lines despite their low absorption, and therefore pointing to a deficiency of cool material if compared with CTTS. The O\textsc{vii} triplet was previously measured for two other WTTS, TWA 5 (Argiroffi et al. 2005) and HD 98800 (Kastner et al. 2004), and in both cases relatively low densities, reminiscent of the coronae of main-sequence stars, were observed.

In conclusion, apart from several T Tauri stars for which the O\textsc{vii} remains undetected, there are so far two clear reports each of high (TW Hya, BP Tau) and of low densities (AB Aur, T Tau) in the cooler plasma component. Although high densities are not a condition to qualify the soft X-ray emission for the accretion scenario, we suggested that the accretion rates and the filling factors would make this scenario unlikely for AB Aur. A specific discussion of the T Tau observation will be given by Güdel et al. (2007b). The triplet line flux ratios do therefore not seem to be reliable indicators for accretion-induced X-rays on T Tau stars.

As for the issue of overabundances of specific elements such as Ne or N, we have indeed found the Ne abundance to be high compared to Fe (4-6 times higher than the solar ratio), but this is the case for all stars except the G-type stars (SU Aur, HP Tau/G2, and HD 283572) and the Herbig star AB Aur. Stars with a high Ne abundance in their X-ray source thus comprise both WTTS and CTTS. A high Ne abundance has also been reported for three members of the TW association, namely TW Hya (Kastner et al. 2002), HD 98800 (Kastner et al. 2004), and TWA 5 (Argiroffi et al. 2005), of which the last two seem to be non-accreting T Tau stars. Therefore, we suggest that the high Ne abundance is not a characteristic property of CTTS, but is common to most young low-mass stars. Studying the abundances of our stellar sample and a sample of other PMS and active MS stars, we have found that the G-type stars on average show abundance ratios of $\langle \text{Fe/Ne} \rangle_G = 1.02$ and $\langle \text{Fe/O} \rangle_G = 2.03$, while we have found significantly lower ratios for K-type stars (on average $\langle \text{Fe/Ne} \rangle_K = 0.22$ and $\langle \text{Fe/O} \rangle_K = 0.58$). It thus seems that the abundance ratios are a function of the spectral type and are similar for PMS stars and more evolved active stars, while we find no trend with respect to accretion. The abundance of N is difficult to measure because the relevant lines are usually rather faint or suppressed by photoelectric absorption. In Chapter 5 we do not find a significant anomaly in AB Aur for which the N\textsc{vii} line could be fitted.

Finally, the outstanding property of the TW Hya spectrum reported previously (Kastner et al. 2002) is the dominance of soft emission relating to cool plasma. Other T Tau stars, however, regularly show hard spectra from which a dominant hot plasma component is inferred (e.g., Preibisch et al. 2005; Güdel et al. 2007a). Here, we have again made specific use of the high-resolution available from the RGS instrument at low photon energies. We have studied the flux ratio of O\textsc{vii}/O\textsc{viii} that is inaccessible to CCD instruments. The O\textsc{vii} line is formed in a relatively narrow temperature range ($\approx 1-4$ MK) with
the emissivity peaking at $\approx 2$ MK, whereas the O\emph{viii} line forms over a wider range centered at somewhat higher temperatures ($\approx 4$ MK). The O\emph{vii}/O\emph{viii} flux ratio is thus a good temperature indicator for the plasma at the cool end of the emission measure distribution. The ratio may be modified by photoelectric absorption, but we have corrected for this effect and determined the single temperature that is equivalent to the unabsorbed flux ratio (see Fig. 3.8). In our sample, it is evident that the accreting stars show an \textit{excess} in the softest emission (expressed by lower $T_{\text{oxy}}$ in Fig. 3.8) that is not present in the non-accreting stars. While this \textit{soft excess} may be of little relevance for the overall emission measure distribution dominated by plasma up to 30 MK, it does systematically alter the lines formed at the lowest temperatures. The excess of cool emission measure could be due to additional volume containing cool plasma, or due to increased densities of low-temperature plasma. We cannot conclusively distinguish between these alternatives, but note that accreting stars with high and low densities as inferred from O\emph{vii} line ratios have now been reported. We conclude that while hot plasma may dominate the X-ray sources of most T Tau stars, a \textit{soft excess} is characteristic of the accreting stars.

What, then, is the mechanism that induces a soft excess in the accreting subsample of our targets? The accretion shock scenario (Kastner et al. 2002) remains a possibility for TW Hya, BP Tau, and V4046 Sgr although the requirements for AB Aur and T Tau are rather demanding, as lower accretion rates than hitherto estimated and/or very large accretion areas on the star are required, in contradiction to the standard magnetic-funnel accretion scenario (Calvet & Gullbring 1998).

Audard et al. (2005) have observed a strongly accreting T Tau star (a so-called EXor object) during an outburst attributed to a strong accretion event. They noted a significant softening of the spectrum during and after outburst, indicating the predominance of cooler plasma, although the X-ray luminosity did not change significantly. They speculated that the accreting material is disrupting the largest magnetic features during outburst, which would also be the hottest if magnetic loops have the same pressure. This would favor emission from smaller, cooler magnetically confined regions in those areas where accretion is active.

Alternatively, the accreting material may not disrupt the magnetic structures but may instead fill them with additional cool material that is not driven into the coronal regions from the chromospheric layers by the mechanism of coronal heating. The plasma in the accreting loops is thus cooler from the outset. Also, the increased electron density increases the cooling efficiency of heated loops, because the cooling losses scale as $n_e^2$. Magnetic loops loaded with accreting material are therefore cooler (Preibisch et al. 2005) if not otherwise heated preferentially.
Chapter 4

X-Ray Emission from T Tauri Stars and the Role of Accretion: Inferences from the XMM-Newton Extended Survey of the Taurus Molecular Cloud

ABSTRACT: T Tau stars display different X-ray properties depending on whether they are accreting (classical T Tau stars; CTTS) or not (weak-line T Tau stars; WTTS). X-ray properties may provide insight into the accretion process between disk and stellar surface. We use data from the XMM-Newton Extended Survey of the Taurus Molecular Cloud (XEST) to study differences in X-ray properties between CTTS and WTTS. We perform correlation and regression analysis between X-ray parameters and stellar properties. We confirm the existence of a X-ray luminosity ($L_X$) vs. mass ($M$) relation, $L_X \propto M^{1.69\pm0.11}$, but this relation is a consequence of X-ray saturation and a mass vs. bolometric luminosity ($L_*$) relation for the TTS with an average age of 2.4 Myr. X-ray saturation indicates $L_X = \text{const}L_*$, although the constant is different for the two subsamples: const $= 10^{-3.73\pm0.05}$ for CTTS and const $= 10^{-3.39\pm0.06}$ for WTTS. Given a similar $L_*$ distribution of both samples, the X-ray luminosity function also reflects a real X-ray deficiency in CTTS, by a factor of $\approx 2$ compared to WTTS. The average electron temperatures $T_{av}$ are correlated with $L_X$ in WTTS but not in CTTS; CTTS sources are on average hotter than WTTS sources. At best marginal dependencies are found between X-ray properties and mass accretion rates or age. The most fundamental properties are the two saturation laws, indicating suppressed $L_X$ for CTTS. We speculate that some of the accreting material in CTTS is cooling active regions to temperatures that may not significantly emit in the X-ray band,
and if they do, high-resolution spectroscopy may be required to identify lines formed in such plasma, while CCD cameras do not detect these components. The similarity of the $L_X$ vs. $T_{\text{av}}$ dependencies in WTTS and main-sequence stars as well as their similar X-ray saturation laws suggests similar physical processes for the hot plasma, i.e., heating and radiation of a magnetic corona.
Optically revealed low mass pre-main sequence stars define the class of T Tauri Stars (TTS). TTS are divided into two families, the classical T Tauri Stars (CTTS) and the weak-line T Tauri Stars (WTTS). CTTS display strong H$\alpha$ lines, a sign that the stars are accreting material from the circumstellar disk, while in WTTS the H$\alpha$ line fluxes are suppressed, a sign that accretion has ceased. Based on infrared observations, Young Stellar Objects (YSO) have instead been ordered in classes according to their infrared (IR) excess. Following this classification, deeply embedded stars at the start of their accretion phase are “Class 0” objects, more evolved protostars still embedded in their envelope are “Class I” objects, stars with a circumstellar disk that show IR excess are “Class II” objects, and stars with no IR excess are “Class III” objects. While the H$\alpha$ classification is based on accretion, the IR excess is a measure of circumstellar material. The Class II objects are dominated by CTTS, while the Class III stars are dominated by WTTS.

Both types of TTS have been found to be strong X-ray emitters. First X-ray detections of individual TTS were made with the Einstein observatory (e.g., Feigelson & DeCampli 1981) and revealed very strong X-ray activity, exceeding the solar level by several orders of magnitude. Many star-forming regions have subsequently been observed with the ROSAT satellite (e.g., Feigelson et al. 1993; Gagné et al. 1995; Neuhauser et al. 1995; Stelzer & Neuhauser 2001), largely increasing the number of X-ray detected TTS. Studies based on H$\alpha$ emission may in fact fail to detect part of the WTTS population, which can easily be identified in X-rays.

The origin of the strong X-ray activity in TTS is not entirely clear. The observed emission in the soft X-ray band above 1 keV is consistent with emission from a scaled-up version of the solar corona. In main-sequence stars X-ray activity is mainly determined by the stellar rotation rate. The activity-rotation relation is given by $L_X/L_\ast \propto P_{\text{rot}}^{-2.6}$ (Gudel et al. 1997a), where $L_X$ is the X-ray luminosity, $L_\ast$ is the stellar photospheric bolometric luminosity, and $P_{\text{rot}}$ is the rotation period of the star. This is consistent with the dynamo mechanism that is present in our Sun, where the magnetic fields are generated through an $\alpha\Omega$-dynamo (Parker 1955). At rotation periods shorter than 2-3 days for G-K stars, the X-ray activity saturates at $\log(L_X/L_\ast) \approx -3$ (Vilhu & Rucinski 1983).

As for pre-main sequence stars, early surveys of the Taurus Molecular Cloud (TMC; Neuhauser et al. 1995; Stelzer & Neuhauser 2001) claimed a rotation-activity relation somewhat similar to the relation for main-sequence stars, but the recent COUP survey of the Orion Nebula Cluster (ONC) found absence of such a relation (Preibisch et al. 2005), suggesting that all stars are in a saturation regime, even for long rotation periods. Young stellar objects, especially
in their early evolutionary stage, are thought to be fully convective, and the generation of magnetic fields through the $\alpha\Omega$-dynamo should not be possible. This suggests that X-rays in low-mass pre-main sequence stars are generated through processes different than in the Sun. New models for X-ray generation through other dynamo concepts have been developed (Kükner & Rüdiger 1999; Giampapa et al. 1996). Alternatively, in CTTS, X-rays could in principle be produced by magnetic star-disk interactions (e.g., Montmerle et al. 2000; Isobe et al. 2003), in accretion shocks (e.g., Lamzin 1999; Kastner et al. 2002; Stelzer & Schmitt 2004), or in shocks at the base of outflows and jets (Güdel et al. 2005a; Kastner et al. 2005).

The influence of a circumstellar disk, and particularly the influence of accretion on X-ray activity is therefore of interest. Former X-ray studies of star forming regions have led to discrepant results. In the Taurus-Auriga complex, Stelzer & Neuhauser (2001) reported higher X-ray luminosities for the non-accreting WTTS stars than for CTTS. In the ONC, Feigelson et al. (2002) concluded from Chandra observations that the presence of circumstellar disks has no influence on the X-ray emission, whereas Flaccomio et al. (2003a), in another Chandra study of the ONC, found $L_X$ and $L_X/L_*$ to be enhanced in WTTS when compared to CTTS. From the recent Chandra Orion Ultradeep Project (COUP), Preibisch et al. (2005) reported the X-ray emission of WTTS to be consistent with the X-ray emission of active Main Sequence (MS) stars, while it is suppressed in CTTS. However, in all these studies the X-ray emission mechanism is consistent with a scaled-up version of a solar corona.

X-ray emission during accretion outbursts has been observed in V1647 Ori (Kastner et al. 2004; Grosso et al. 2005; Kastner et al. 2006) and in V1118 Ori (Audard et al. 2005). The X-ray luminosity increased by a factor of 50 during the outburst in V1647 Ori, and the spectrum hardened. On the other hand, the X-ray luminosity of V1118 Ori remained at the same level as during the pre-outburst phase, while the spectrum became softer.

Possible signs of accretion-induced X-ray emission are revealed in a few high-resolution spectra of CTTS. High electron densities were measured in the spectra of TW Hya (Kastner et al. 2002; Stelzer & Schmitt 2004), BP Tau (Schmitt et al. 2005; Robrade & Schmitt 2006), and V4046 Sgr (Günther et al. 2006), and were interpreted as indications of X-ray production in accretion shocks. Other spectroscopic features that also suggested an accretion shock scenario are the low electron temperature dominating the plasma in TW Hya (a few MK, as expected from shock-induced heating) and abundance anomalies. Stelzer & Schmitt (2004) interpreted the high Ne/Fe abundance ratio as being due to Fe depletion by condensation into grain in the accretion disk. Drake et al. (2005b) reported a substantially larger Ne/O ratio in the spectrum of TW Hya than in the spectra of the other studied stars, and they proposed to use this ratio as a diagnostic for metal depletion in the circumstellar disk of accreting stars.
Work on high-resolution X-ray spectroscopy was subsequently extended (Chapter 3) to a sample of 9 pre-main sequence stars with different accretion properties. The main result of that work is the identification of an excess of cool plasma measured in the accreting stars, when compared to WTTS. The origin of this soft excess is unclear. Further evidence for a strong soft excess in the CTTS is revealed in the extraordinary X-ray spectrum of T Tau (Güdel et al. 2007b). In this case, however, the electron density (derived from spectral lines formed at low temperatures) is low, \( n_e \lesssim 10^{10} \text{ cm}^{-3} \). The density, in case of accretion shocks, can be estimated using the strong shock condition \( n_2 = 4n_1 \), where \( n_1 \) and \( n_2 \) are the pre-shock and post-shock densities, respectively. The density \( n_1 \) can be derived from the accretion mass rate and the accreting area on the stellar surface: \( \dot{M} \approx 4\pi R^2 f v_{ff} n_e m_p \), where \( f \) is the surface filling factor of the accretion flow, and \( v_{ff} = (2GM/R)^{1/2} \) is the free-fall velocity. Using the accretion rate \( \dot{M} \approx (3 - 6) \times 10^{-8} M_\odot \text{ yr}^{-1} \) for T Tau (White & Ghez 2001; Calvet et al. 2004) we obtain \( n_2 = (1.1 - 2.2) \times 10^{11}/f \) (Güdel et al. 2007b). Even in the extreme case that \( f = 10\% \), we expect a density \( \gtrsim 10^{12} \text{ cm}^{-3} \), i.e. orders of magnitude higher than the measured value.

The aim of the this chapter is to study the role of accretion in the overall X-ray properties of pre-main sequence stars in the Taurus-Auriga Molecular Cloud, by coherently comparing samples of CTTS and WTTS. Our analysis is complementary to the COUP survey work, and we will present our results along largely similar lines (see Preibisch et al. 2005 for COUP). Indeed one of the main purposes of the present work is a qualitative comparison of the Taurus results with those obtained from the Orion sample. We do not, however, present issues related to rotation; rotation-activity relations will be separately discussed in a dedicated paper (Briggs et al. 2007).

This chapter is structured as follows. In Sect. 4.2 we describe the stellar sample used in this work, and in Sect. 4.3 we summarize the relevant steps of the data reduction. We present our results in Sect. 4.4, and discuss them in Sect. 4.5. We summarize our results and conclude in Sect. 4.6.

4.2 Studying X-Rays in the Taurus Molecular Cloud

4.2.1 The Taurus Molecular Cloud

We will address questions on X-ray production in accreting and non-accreting T Tauri stars using data from the XMM-Newton Extended Survey of the Taurus Molecular Cloud (XEST, Güdel et al. 2007a).

The Taurus Molecular Cloud (TMC) varies in significant ways from the Orion Nebula Cluster and makes our study an important complement to the COUP survey. The TMC has, as the nearest large, star-forming region (dis-
Chapter 4. X-Ray Emission from T Tauri Stars and the Role of Accretion in XEST

tance ≈ 140 pc, Loinard et al. 2005), played a fundamental role in our understanding of low-mass star formation. It features several loosely associated but otherwise rather isolated molecular cores, each of which produces one or only a few low-mass stars, different from the much denser cores in ρ Oph or in Orion. TMC shows a low stellar density of only 1–10 stars pc$^{-2}$ (e.g., Gómez et al. 1993). In contrast to the very dense environment in the Orion Nebula Cluster, strong mutual influence due to outflows, jets, or gravitational effects is therefore minimized. Also, strong winds and UV radiation fields of OB stars are present in Orion but absent in the TMC.

The TMC has provided the best-characterized sample of CTTS and WTTS, many of which have been subject to detailed studies; see, e.g., the seminal work by Kenyon & Hartmann (1995) that concerns, among other things, the evolutionary history of T Tau stars and their disk+envelope environment, mostly based on optical and infrared observations.

It is therefore little surprising that comprehensive X-ray studies of selected objects as well as surveys have been performed with several previous X-ray satellites; for X-ray survey work see, e.g., the papers by Feigelson et al. (1987), Walter et al. (1988), Bouvier (1990), Strom et al. (1990), Strom & Strom (1994), Damiani & Micela (1995), Damiani et al. (1995), Neuhäuser et al. (1995), and Stelzer & Neuhäuser (2001). Issues we are studying in this chapter have variously been studied in these surveys before, although, as argued by Güdel et al. (2007a) and below, the present XEST project is more sensitive and provides us with a near-complete sample of X-ray detected TTS in the surveyed area, thus minimizing selection and detection bias.

4.2.2 The XEST Sample of T Tau Stars

The XEST project is an X-ray study of the most populated regions (comprising an area of ≈ 5 square degrees) of the Taurus Molecular Cloud. The survey consists of 28 XMM-Newton exposures. The 19 initial observations of the project (of approximately 30 ks duration each, see Table 1 of Güdel et al. 2007a) were complemented by 9 exposures from other projects or from the archive. Also, 6 Chandra observations have been used in XEST, to add information on a few sources not detected with XMM-Newton, or binary information (see Güdel et al. 2007a).

To distinguish between CTTS and WTTS we use the classification given in col. 10 of Table 11 of Güdel et al. (2007a). This classification is substantially based on the equivalent width of the Hα line (EW[Hα]). For spectral types G and K, stars with EW(Hα) ≥ 5 Å are defined as CTTS, while other stars are defined as WTTS. For early-M spectral types, the boundary between CTTS and WTTS was set at EW(Hα) = 10 Å and for mid-M spectral types at EW(Hα) = 20 Å. Stars with late-M spectral type are mainly Brown Dwarfs (BDs), and given their low optical continuum, a clear accretion criterion is
4.2. Studying X-Rays in the Taurus Molecular Cloud

difficult to provide. For this reason, BDs were treated as a class of their own and are not used in our comparison studies of accretors vs. non-accretors (but were included in the “total” samples when appropriate). For further details, see Güdel et al. (2007a) and Grosso et al. (2007a). YSO IR types were used to classify borderline cases and protostars. In summary, protostars have been classified as type 0 or 1 (Class 0 and I, respectively), CTTS are type 2 objects, WTTS are type 3 objects, and BDs are classified as type 4. Type 5 is assigned to Herbig Ae/Be stars, while stars with uncertain classification are assigned to type 9. We will use these designations in our illustrations below.

We emphasize the near-completeness of XEST with regard to X-ray detections of TTS. Güdel et al. (2007a) provide the detection statistics (their Table 12): A total of 126 out of the 159 TMC members surveyed with XMM-Newton were detected in X-rays. Among these are 55 detected CTTS and 49 detected WTTS (out of the 65 and 50 surveyed targets), corresponding to a detection fraction of 85% and 98%, respectively. Almost all objects have been found comfortably above the approximate detection limit of \( L_X \approx 10^{28} \text{ erg s}^{-1} \), indicating that TTS generally emit at levels between \( 10^{29} - 10^{31} \text{ erg s}^{-1} \), exceptions being lowest-mass stars and brown dwarfs. Most of the non-detected objects have been recognized as stars that are strongly absorbed (e.g., by their own disks) or as stars of very low mass (Güdel et al. 2007a). XEST is the first X-ray survey of TMC that reaches completeness fractions near unity, and therefore minimizes detection bias and unknown effects of upper limits to correlation studies as performed here. It provides, in this regard, an ideal comparison with the COUP results (Preibisch et al. 2005).

A few sources were excluded from consideration in the present work. These are the four stars that show composite X-ray spectra possibly originating from two different sources (DG Tau A, GV Tau, DP Tau, and CW Tau; Güdel et al. 2007c), and three stars which show a decreasing light curve throughout the observation (DH Tau, FS Tau AC, and V830 Tau); these light curves probably describe the late phases of large flares. Further, the deeply embedded protostar L1551 IRS5, which shows lightly absorbed X-ray emission that may be attributed to the jets (Favata et al. 2002; Bally et al. 2003), was also excluded, and so were the two Herbig stars (AB Aur, V892 Tau). In some correlation studies, we do consider objects for which upper limits to \( L_X \) have been estimated in Güdel et al. (2007a), but will not consider non-detections without such estimate (as, for example, if the absorption is unknown).

Our final, basic sample of TTS then consists of 56 CTTS and 49 WTTS. Among the X-ray detections, there are also 8 protostars, 8 BDs, 2 Herbig stars, and 4 stars with uncertain classification. Smaller subsamples may be used if parameters of interest were not available.

When \( L_\ast \) is involved in a correlation, we excluded all stars that are apparently located below the Zero-Age Main Sequence (ZAMS) in the Hertzsprung-Russell diagram (Fig. 10 of Güdel et al. 2007a). These stars are HH 30, IRAS
S04301+261, Haro 6-5 B, HBC 353, and HBC 352. Their location in the HRD is likely to be due to inaccurate photometry.

Many of the stellar counterparts to our X-ray sources are unresolved binaries or multiples. In total, 45 out of the 159 stellar systems surveyed by *XMM-Newton* are multiple (Güdel et al. 2007a). If - as we will find below in general, and as has been reported in earlier studies of T Tauri stars (e.g., Preibisch et al. 2005) - $L_X$ scales with $L_*$, then this also holds for the sum of the $L_X$ with respect to the sum of the component $L_*$. Binarity does therefore not influence comparisons between $L_X$ and $L_*$. When correlating $L_X$ with stellar mass, we will find that more massive stars are in general brighter. In the case of binaries, the more massive component (usually the more luminous “primary” star) will thus dominate the X-ray emission. We have used the primary mass for the stellar systems if available; we therefore expect the influence of the companions on our correlations to be small. We will also present tests with the subsample of single stars below.

### 4.3 Data Reduction and Analysis

The XEST survey is principally based on CCD camera exposures, but is complemented with high-resolution grating spectra for a few bright stars (Chapter 3), and with Optical Monitor observations (Audard et al. 2007). The three EPICs on board *XMM-Newton* are CCD-based X-ray cameras that collect photons from the three telescopes. Two EPIC detectors are of the MOS type (Turner et al. 2001) and one is of the PN type (Strüder et al. 2001). They are sensitive in the energy range of 0.15-15 keV with a spectral resolving power of $E/\Delta E = 20-50$.

The data were reduced using the Science Analysis System (SAS) version 6.1. A detailed description of all data reduction procedures is given in Sect. 4 of Güdel et al. (2007a).

Source and background spectra have been obtained for each instrument using data during the Good Time Intervals (GTIs, i.e., intervals that do not include flaring background). Further, time intervals with obvious strong stellar flares were also excluded from the spectra in order to avoid bias of our results by episodically heated very hot plasma.

One PN spectrum usually provides more counts than the two MOS spectra together. We therefore only used the PN data for the spectral analysis, except for the sources for which PN data were not available (e.g., because the PN was not operational, or the sources fell into a PN CCD gap).

The spectral fits were performed using two different approaches in the full energy band. First, we have fitted the spectra using a conventional one- or two-component spectral model (1-$T$ and 2-$T$), both components being subject to a common photoelectric absorption. In this approach, the hydrogen column
density, $N_H$, two temperatures ($T_{1,2}$) and two emission measures (EM$_{1,2}$) are fitted in XSPEC (Arnaud 1996) using the vapec thermal collisional-ionization equilibrium model.

In the second approach, the spectra were fitted with a model consisting of a continuous emission measure distribution (EMD) as found for pre-main sequence and active ZAMS stars (Argiroffi et al. 2004; García-Alvarez et al. 2005; Scelsi et al. 2005; Chapter 2). This model has been described in Chapter 1 (Sect. 1.6). It consists of a grid of 20 thermal components binned to intervals of $d \log T = 0.1$ from $\log T = 6$ to $\log T = 7.9$, arranged such that they form an EMD with a peak at a temperature $T_0$ and two power-laws toward lower and higher temperatures with power-law indices $\alpha$ and $\beta$, respectively. Given the poor sensitivity of CCD spectra at low temperatures, $\alpha$ was kept fixed at 2, consistent with values found in previous studies (Argiroffi et al. 2004; Chapter 2), while we let $\beta$ free to vary (between $-3 \leq \beta \leq 1$). The absorbing hydrogen column density $N_H$ was also fitted to the data. The abundances were fixed at values typical for pre-main sequence stars or very active zero-age main-sequence stars (Argiroffi et al. 2004; García-Alvarez et al. 2005; Scelsi et al. 2005, Chapter 2)$^1$. For further details, see Güdel et al. (2007a).

For each star and each model we computed the average temperature ($T_{av}$) as the logarithmic average of all temperatures used in the fit, applying the emission measures as weights. The X-ray luminosity ($L_X$) was computed in the energy range 0.3-10 keV from the best-fit model assuming a distance of 140 pc.

Of the 126 members detected in XEST, 22 were detected in two different exposures. In those cases, two separate spectral fits were made. For correlations of X-ray parameters with stellar properties, we used logarithmic averages of the results from the two fits. On the other hand, if we correlate X-ray properties with each other, we treat the two spectral fit results from the same source as different entries.

Results from the spectral fits are given in Table 5 (for the EMD fits) and Table 6 (for the 1-T and 2-T fits) of Güdel et al. (2007a). We use the results from the EMD interpretation to perform statistical correlations below.

### 4.4 Results

Motivated by results from previous X-ray studies and in particular guided by the COUP work (Preibisch et al. 2005), we now seek systematics in the X-ray emission by correlating X-ray parameters first with fundamental stellar parameters, and then also seeking correlations among the X-ray parameters.

---

$^1$The abundance values used are, with respect to the solar photospheric abundances of Anders & Grevesse (1989): C=0.45, N=0.788, O=0.426, Ne=0.832, Mg=0.263, Al=0.5, Si=0.309, S=0.417, Ar=0.55, Ca=0.195, Fe=0.195, Ni=0.195
themselves. We will consider the fundamental stellar properties of mass and bolometric luminosity, accretion rate, and age, but we will not discuss rotation properties here (see Briggs et al. 2007 for a detailed study). The basic X-ray properties used for our correlations are the X-ray luminosity $L_X$ in the 0.3-10 keV band and the average electron temperature $T_{av}$. One of the main goals of this section is to seek differences between CTTS and WTTS. We will compare our findings with those of COUP and some other previous work in Sect. 4.5.

### 4.4.1 Correlations Between X-Rays and Stellar Parameters

**Correlation with Mass**

In Fig. 4.1 we plot the X-ray luminosity, $L_X$ (in erg s$^{-1}$), as a function of the stellar mass ($M$, in units of the solar mass, $M_\odot$, from Table 10 in Güdel et al. 2007a). In the left panel we show the relation for all types of objects in our sample. Different symbols are used to mark different object types (see panel in the figure). Upper limits for non-detections are marked with arrows. In the middle and right panels the same relation is shown separately for CTTS and WTTS, respectively. A clear correlation is found between the two parameters in all three plots, in the sense that $L_X$ increases with mass. The correlation coefficients are 0.79 for the whole sample (99 entries), 0.74 for the CTTS (45 entries), and 0.84 for the WTTS (43 entries). We computed the significance of the correlation using correlation tests in ASURV (LaValley et al. 1992; specifically, the Cox hazard model, Kendall’s tau, and Spearman’s rho have been used) and found a probability of $< 0.01\%$ that the parameters are uncorrelated.
4.4. Results

in each of the three cases. As for all subsequent statistical correlation studies, we summarize these parameters in Table 4.1.

We computed linear regression functions for the logarithms of the two parameters, of the form $\log y = a + b \log x$, using the parametric estimation maximization (EM) algorithm in ASURV, which implements the methods presented by Isobe et al. (1986). We find the regression functions $\log L_X = (1.69 \pm 0.11) \log M + (30.33 \pm 0.06)$ for the full stellar sample, $\log L_X = (1.70 \pm 0.20) \log M + (30.13 \pm 0.09)$ for the CTTS, and $\log L_X = (1.78 \pm 0.17) \log M + (30.57 \pm 0.09)$ for the WTTS. The regression parameters are also listed in Table 4.1, as for all subsequent regression analyses. In the ONC sample, Preibisch et al. (2005) found the similar linear regression $\log L_X = (1.44 \pm 0.10) \log M + (30.37 \pm 0.06)$ for all stars with masses $< 2 M_\odot$ using the same algorithm.

The EM algorithm is an ordinary least-square (OLS) regression of the dependent variable $y$ ($L_X$ in this case) against the independent variable $x$ ($M$). When using this method, we assume that $L_X$ is functionally dependent on the given mass (Isobe et al. 1990). However, the $M$ values are also uncertain, and assuming a functional dependence a priori may not be correct. We therefore also computed the linear regression using the bisector OLS method after Isobe et al. (1990), which treats the variables symmetrically. In this case, we find $\log L_X = (1.91 \pm 0.11) \log M + (30.44 \pm 0.05)$ for all stars together, $\log L_X = (1.98 \pm 0.20) \log M + (30.24 \pm 0.06)$ for the CTTS, and $\log L_X = (2.08 \pm 0.17) \log M + (30.69 \pm 0.07)$ for the WTTS (see Table 4.1). The slopes for the bisector OLS are slightly steeper than in the EM algorithm. However, the values for CTTS and WTTS agree within one sigma, and we caution that the upper limits for non-detections are not taken into account in the bisector linear regression method.

We verified this trend for the subsample of stars that have not been recognized as multiples. We find the regression lines $\log L_X = (1.72 \pm 0.12) \log M + (30.39 \pm 0.07)$ using the EM algorithm, and $\log L_X = (1.85 \pm 0.11) \log M + (30.48 \pm 0.07)$ using the bisector algorithm. These results are fully consistent with the results for the total sample.

Differences are present between the CTTS and WTTS stellar samples. While the slopes found in the correlations for CTTS and WTTS are consistent within 1$\sigma$, the intercept of WTTS at $1 M_\odot$ ($\log M/M_\odot = 0$) is $\approx 0.45$ dex larger than the intercept for CTTS. This lets us anticipate a larger average $L_X$ in WTTS. The correlation is better determined for WTTS, as judged from a slightly higher correlation coefficient and a smaller error in the slope. Furthermore, the standard deviation, $\sigma$, of the points with respect to the regression function from the EM algorithm is slightly larger for CTTS (0.45) than for WTTS (0.38).
Table 4.1: Summary of results found for the different correlations. In the third column, n is the number of stars used in the statistic. $P$ is the probability that the parameters are uncorrelated (computed with ASURV), $C$ is the correlation coefficient and $\sigma$ is the standard deviation from the EM algorithm. The intercept of the linear regression is $a$, and $b$ is the slope. Errors are 1-sigma rms values for the respective variables.

<table>
<thead>
<tr>
<th>Correlation</th>
<th>stellar sample</th>
<th>n</th>
<th>EM algorithm $a$</th>
<th>EM algorithm $b$</th>
<th>bisector algorithm $a$</th>
<th>bisector algorithm $b$</th>
<th>$P$</th>
<th>$C$</th>
<th>$\sigma$</th>
</tr>
</thead>
<tbody>
<tr>
<td>$L_X$ vs $M$</td>
<td>all</td>
<td>99</td>
<td>$30.33 \pm 0.06$</td>
<td>$1.69 \pm 0.11$</td>
<td>$30.44 \pm 0.05$</td>
<td>$1.91 \pm 0.11$</td>
<td>&lt; 0.01%</td>
<td>0.79</td>
<td>0.45</td>
</tr>
<tr>
<td>$L_X$ vs $M$</td>
<td>CTTS</td>
<td>45</td>
<td>$30.13 \pm 0.09$</td>
<td>$1.70 \pm 0.20$</td>
<td>$30.24 \pm 0.06$</td>
<td>$1.98 \pm 0.20$</td>
<td>&lt; 0.01%</td>
<td>0.74</td>
<td>0.45</td>
</tr>
<tr>
<td>$L_X$ vs $M$</td>
<td>WTTS</td>
<td>43</td>
<td>$30.57 \pm 0.09$</td>
<td>$1.78 \pm 0.17$</td>
<td>$30.69 \pm 0.07$</td>
<td>$2.08 \pm 0.17$</td>
<td>&lt; 0.01%</td>
<td>0.84</td>
<td>0.38</td>
</tr>
<tr>
<td>$T_{av}$ vs $L_X$</td>
<td>CTTS</td>
<td>19</td>
<td>$6.45 \pm 2.31$</td>
<td>$0.01 \pm 0.08$</td>
<td>$-17.95 \pm 13.25$</td>
<td>$0.85 \pm 0.45$</td>
<td>43-80%</td>
<td>0.06</td>
<td>0.22</td>
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<tr>
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<td>$2.53 \pm 0.81$</td>
<td>$0.15 \pm 0.03$</td>
<td>$0.13 \pm 0.93$</td>
<td>$0.23 \pm 0.03$</td>
<td>&lt; 0.01%</td>
<td>0.69</td>
<td>0.12</td>
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<td>$T_{av}$ vs $F_X$</td>
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<td>$6.77 \pm 0.76$</td>
<td>$0.05 \pm 0.12$</td>
<td>$1.67 \pm 1.11$</td>
<td>$0.86 \pm 0.18$</td>
<td>62-42%</td>
<td>0.11</td>
<td>0.22</td>
</tr>
<tr>
<td>$T_{av}$ vs $F_X$</td>
<td>WTTS</td>
<td>32</td>
<td>$5.75 \pm 0.20$</td>
<td>$0.18 \pm 0.03$</td>
<td>$5.21 \pm 0.24$</td>
<td>$0.26 \pm 0.03$</td>
<td>&lt; 0.01%</td>
<td>0.72</td>
<td>0.11</td>
</tr>
<tr>
<td>$L_X$ vs $L_\ast /L_\odot$</td>
<td>all</td>
<td>108</td>
<td>$30.00 \pm 0.05$</td>
<td>$1.05 \pm 0.06$</td>
<td>$30.07 \pm 0.04$</td>
<td>$1.11 \pm 0.05$</td>
<td>&lt; 0.01%</td>
<td>0.83</td>
<td>0.44</td>
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<tr>
<td>$L_X$ vs $L_\ast /L_\odot$</td>
<td>CTTS</td>
<td>48</td>
<td>$29.83 \pm 0.06$</td>
<td>$1.16 \pm 0.09$</td>
<td>$29.89 \pm 0.05$</td>
<td>$1.20 \pm 0.10$</td>
<td>&lt; 0.01%</td>
<td>0.84</td>
<td>0.39</td>
</tr>
<tr>
<td>$L_X$ vs $L_\ast /L_\odot$</td>
<td>WTTS</td>
<td>44</td>
<td>$30.22 \pm 0.08$</td>
<td>$1.06 \pm 0.10$</td>
<td>$30.31 \pm 0.06$</td>
<td>$1.25 \pm 0.09$</td>
<td>&lt; 0.01%</td>
<td>0.85</td>
<td>0.41</td>
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<td>$L_X/L_X(M)$ vs $M$</td>
<td>CTTS</td>
<td>37</td>
<td>$-4.05 \pm 1.19$</td>
<td>$-0.48 \pm 0.15$</td>
<td>$-8.32 \pm 1.11$</td>
<td>$-1.02 \pm 0.14$</td>
<td>$0.14 - 0.52$</td>
<td>-0.47</td>
<td>0.57</td>
</tr>
<tr>
<td>$L_X(M = 1M_\odot)$ vs age</td>
<td>all</td>
<td>93</td>
<td>$30.45 \pm 0.06$</td>
<td>$-0.36 \pm 0.11$</td>
<td>$30.69 \pm 0.06$</td>
<td>$-1.02 \pm 0.07$</td>
<td>$0.10 - 0.38$</td>
<td>-0.31</td>
<td>0.44</td>
</tr>
<tr>
<td>$L_\ast /L_\odot$ vs $M$</td>
<td>all</td>
<td>113</td>
<td>$0.23 \pm 0.04$</td>
<td>$1.49 \pm 0.07$</td>
<td>$0.93 \pm 0.04$</td>
<td>$1.65 \pm 0.06$</td>
<td>&lt; 0.01%</td>
<td>0.90</td>
<td>0.31</td>
</tr>
</tbody>
</table>
4.4. Results

Figure 4.2: Evolution of the normalized X-ray emission. $L_X$ has been normalized with the predicted values from the $L_X$-mass relation (see text for details). The linear regression computed with the EM algorithm is plotted (solid line) together with errors in the slopes (dashed line). Symbols mark different types of stars.

**Evolution of X-Ray Emission**

Here, we discuss the evolution of the X-ray emission with age. Among main-sequence (MS) stars, $L_X$ is correlated with rotation and anti-correlated with age. The common explanation is that magnetic activity is directly related to the stellar rotation, and the latter decays with age because of magnetic braking. However, TMC PMS stars do not show the relation between $L_X$ and rotation observed for MS stars (Briggs et al. 2007). On the other hand, $L_X$ decreases during the evolution of pre-main sequence stars, provided that a common X-ray saturation law applies (see below), because $L_\ast$ decreases along the Hayashi track.

We have found (see Sect. 4.4.1) that $L_X$ shows a strong correlation with mass. In order to avoid an interrelationship between the correlations, we normalize the measured $L_X$ with $L_X(M)$ predicted by the correlation with mass ($L_X(M) = 10^{30.33} M^{1.69} \text{ erg s}^{-1}$) and multiply with the $L_X$ expected for a 1 $M_\odot$ star ($10^{30.33} \text{ erg s}^{-1}$). We designate this quantity by $L_X(M = 1M_\odot)$. In Fig. 4.2 we plot $L_X(M = 1M_\odot)$ as a function of age. A slight decline in $L_X$ is found between 0.1 and 10 Myr. The correlation coefficient is $C = -0.31$ for 93 entries. The tests in ASURV give probabilities between 0.1% and 0.38% that age and $L_X(M = 1M_\odot)$ are uncorrelated. We
have computed a linear regression with the EM algorithm in ASURV and find 
\[ \log\left( \frac{L_X}{L_X(M)} \right) = (-0.36 \pm 0.11) \log \text{(age)} + (30.45 \pm 0.06) \] (where age is in Myr). Further, we have tested the linear regression and the correlation probability when we neglect the two youngest stars (V410 X4 and LkH\alpha 358) and found the linear regression to be consistent within error bars with the above relation, with a probability of \( P < 1\% \) for no correlation. Further, as a test, we have computed the linear regression using the bisector algorithm. We find a much steeper slope of \(-1.02 \pm 0.07\) (Table 4.1), indicating that the linear regression is nevertheless only marginal, and the scatter is dominated by other contributions.

Preibisch & Feigelson (2005) reported correlations consistent with ours, applying the EM algorithm to the ONC data. They correlated \( L_X \) with age in mass-stratified subsamples of the surveyed stars and found \( L_X \) to decrease with age with slopes ranging from -0.2 to -0.5, i.e. fully consistent with the slope found in Fig. 4.2 (mass-stratified analysis for our sample also indicates decreasing \( L_X \) in some mass bins but not in others; our statistics are too poor for this purpose).

**Mass Accretion Rates**

We now directly compare the X-ray parameters derived from our spectral fits \( (L_X, L_X/L_*, \text{ and } T_{av}) \) with the previously determined mass accretion rates \( (\dot{M}, \text{ in } M_\odot \text{ yr}^{-1}) \). We use \( \dot{M} \) listed in the XEST catalog (Güdel et al. 2007a, and references therein). Accretion rates may be variable, and various methods for their determination may produce somewhat different results. If different values were found for a given star in the literature, the range of \( \dot{M} \) is marked by a horizontal line in our figures.

When comparing \( L_X \) with \( \dot{M} \), some caution is in order. In Sect. 4.4.1 we have shown that a tight relation exists between \( L_X \) and the stellar mass. Further, a clear relation between \( \dot{M} \) and the stellar mass has been found for Class II objects in the literature (e.g., Muzerolle et al. 2003, 2005; Calvet et al. 2004). Combining the \( L_X \)-mass and \( M \)-mass relations, we expect \( L_X \) to correlate with \( M \) as well. However, here we are interested in testing if an intrinsic relation between the latter two parameters exists that is not a consequence of the two former relations. Calvet et al. (2004) have used evolutionary tracks of Siess et al. (2000) (consistent with the XEST survey) to find a relation \( \dot{M} \propto M^{1.95} \) in the mass range between 0.02 and 3 \( M_\odot \). Similarly, Muzerolle et al. (2003) and Muzerolle et al. (2005) found \( \dot{M} \propto M^2 \) and \( \dot{M} \propto M^{2.1} \) respectively, using different evolutionary tracks. We therefore adopt the relation \( \log \dot{M} \approx 2 \log M - 7.5 \). Further, we use \( \log L_X = 1.69 \log M + 30.33 \) (Sect. 4.4.1), and we then compute the expected \( L_X \) for each \( \dot{M} \) value, namely \( \log L_X(\dot{M}) = 0.85 \log \dot{M} + 36.67 \).

In Fig. 4.3 we plot the ratio of \( L_X/L_X(\dot{M}) \) as a function of \( \dot{M} \) for class 2 objects. We will expect that the values scatter around a constant if this
4.4. Results

Figure 4.3: Residual X-ray luminosity for CTTS (after normalization with the $M - L_X$ and $M - \dot{M}$ relation) as a function of the mass accretion rate. Regression lines obtained using the EM algorithm (red) and the bisector algorithm (blue) are plotted with their respective errors in the slope (dashed lines). The regression lines are computed only using the stars plotted with black circles, while the two stars with $\dot{M}$ smaller than $10^{-10} \, M_\odot \, yr^{-1}$ were ignored (see text for more details).

ratio were determined only by the $M - L_X$ and $M - \dot{M}$ relations. We find a very large scatter for any given $\dot{M}$ (2–3 orders of magnitude) but, using a regression analysis, a tendency for weak accretors to show higher $L_X$, compared to strong accretors. However, if we exclude the two stars with the smallest accretion rates (plotted with blue symbols in Fig. 4.3), the correlation is less clear. In this case, the correlation coefficient is $C = -0.47$ for 37 data points. Nevertheless, the probability, computed in ASURV for the EM algorithm, that no correlation is present is only $P = 0.1\% - 0.5\%$. We have computed the linear regression using different methods. With the EM algorithm we find $\log L_X/L_X(\dot{M}) = (-0.48 \pm 0.15) \log \dot{M} - (4.05 \pm 1.19)$ (Table 4.1). Using the bisector algorithm, however, the slope is found to be $-1.02 \pm 0.14$, i.e., more than 3 sigma steeper than the slope found with the EM algorithm. The entries for two stars with low $\dot{M}$ (plotted in blue) are consistent with the linear regression found with the EM algorithm. We conclude that the two parameters are not evenly distributed, but that a linear regression of the logarithmic values cannot clearly be claimed.

In Fig. 4.4 we plot $L_X/L_*$ and $T_{\text{av}}$ as a function of the accretion rate. Arrows represent upper limits for $\dot{M}$. In both cases no correlation is evident.
Figure 4.4: Left panel: Fractional X-ray luminosity vs. $\dot{M}$. Right panel: $T_{\text{av}}$ vs. $\dot{M}$. Different symbol sizes represent different object types as defined in the figures. Arrows represent upper limits for the accretion rates.

**Correlation with Bolometric Luminosity**

In Fig. 4.5 we plot $L_X$ as a function of the stellar bolometric luminosity $L_*$ (from Güdel et al. 2007a and references therein). The lines corresponding to $L_X/L_* = 10^{-3}$, $10^{-4}$, and $10^{-5}$ are also shown. In the left panel, all stars are plotted, with different symbols for each stellar class as described in the figure. We again excluded from the plot the stars mentioned in Sect. 4.2. Upper limits for non-detections are marked with arrows. By far most of the stars are located between $L_X/L_* = 10^{-3}$ and $L_X/L_* = 10^{-4}$. In the middle and right panels, we present CTTS and WTTS separately.

The correlation coefficients are 0.83 for the full stellar sample (108 entries), and 0.84 (48 entries) and 0.85 (44 entries) for CTTS and WTTS, respectively. Probabilities for the absence of a correlation are very small, $P < 0.01\%$. We computed linear regression lines with the EM algorithm in ASURV. For the full sample, we find $\log L_X = (1.05 \pm 0.06) \log L_*/L_\odot + (30.00 \pm 0.05)$, whereas for CTTS and WTTS, $\log L_X = (1.16 \pm 0.09) \log L_*/L_\odot + (29.83 \pm 0.06)$ and $\log L_X = (1.06 \pm 0.10) \log L_*/L_\odot + (30.22 \pm 0.08)$, respectively. Fig. 4.5c shows one WTTS at $L_* / L_\odot \approx 0.01$ with a rather high $L_X \approx 5 \times 10^{29}$ erg s$^{-1}$ (KPNO-Tau 8 = XEST-09-022). Not considering this object, the slope of the regression slightly steepens to $1.17 \pm 0.09$, which is only marginally different from the slope based on all WTTS. The standard deviation at the same time marginally decreases from 0.41 to 0.36.

Again, we also computed a linear regression using the bisector OLS algorithm that treats both $L_*$ and $L_X$ as independent variables. The slopes are very similar to the those reported above (Table 4.1). The important distinction between CTTS and WTTS is that the latter clearly tend to be located at higher $L_X/L_*$ (see also below): at $L_* / L_\odot = 1$ the CTTS show an average
4.4. Results

Figure 4.5: $L_X$ as a function of the $L_\star$. Left: for all stars; the different symbols describe different classes of stars, while arrows are upper limits for non-detections. Middle: same for CTTS (type 2). Right: same for WTTS (type 3). The horizontal bars show the ranges of literature values for $L_\star$.

$\log L_X = 29.83$ [erg s$^{-1}$], while for the WTTS, $\log L_X = 30.22$ [erg s$^{-1}$].

The regressions are thus compatible with a linear relation between $L_X$ and $L_\star$, and therefore $L_X/L_\star$ is, on average for a given $L_\star$, a constant between $10^{-4}$ and $10^{-3}$ regardless of $L_\star$. This is reminiscent of the situation among very active, rapidly rotating main-sequence or evolved subgiant stars that saturate at fractional X-ray luminosities of the same order, provided they rotate sufficiently rapidly. We thus find that the majority of our TTS are in a saturated state. A consequence of this would be that rotation no longer controls the X-ray output, as suggested by Preibisch et al. (2005) for the Orion sample. This is discussed for the XEST sample by Briggs et al. (2007). Below, we will specifically study whether the $L_X/L_\star$ relation is different for CTTS and for WTTS.

The correlation found for the full stellar sample is compatible with the relation found in Orion: $\log L_X = (1.04 \pm 0.06) \log(L_\star/L_\odot) + (30.00 \pm 0.04)$ (Preibisch et al. 2005). The slope found for the WTTS in the ONC is also consistent with our results within the error bars. For CTTS, on the other hand, Preibisch et al. (2005) found a very large scatter in the correlation. This is not observed in our XEST sample; we rather see similar scatter for CTTS and WTTS, as demonstrated by the similar correlation coefficients, the similar errors in the slope, and the similar standard deviations. Preibisch et al. (2005) suggested that strong accretion could lead to larger errors in the determination of stellar luminosity and the effective temperature.

**Fractional X-ray Luminosity $L_X/L_\star$**

In Fig. 4.6 we plot the histogram for the distribution of $\log(L_X/L_\star)$ for CTTS (grey) and WTTS (white). The two populations are different, with WTTS
having a larger mean $\log(L_X/L_\odot)$. We fitted each of the two histograms with a Gaussian function and computed the mean and its errors. For CTTS, we find $\langle \log(L_X/L_\odot) \rangle = -3.73 \pm 0.05$, while for WTTS, $\langle \log(L_X/L_\odot) \rangle = -3.39 \pm 0.06$.

A more rigorous test is based on the Kaplan-Meier estimators as computed in ASURV, which implements the methods presented by Feigelson & Nelson (1985). This method also accounts for the upper limits in $L_X$ for the non-detections. The results are plotted in Fig. 4.7. The solid line represents the CTTS, the dotted line the WTTS. The WTTS distribution is shifted toward larger $\log(L_X/L_\odot)$ compared to the CTTS distribution by a factor of approximately 2. We find $\langle \log(L_X/L_\odot) \rangle = -3.72 \pm 0.06$ and $\langle \log(L_X/L_\odot) \rangle = -3.36 \pm 0.07$ for CTTS and WTTS, respectively, in full agreement with the Gaussian fit. Judged from a two-sample test based on the Wilcoxon test and logrank test in ASURV, the probability that the two distributions are obtained from the same parent population is very low, namely $P = 0.01\%-0.03\%$.

Again, we test this result using the subsample of stars that have not been recognized as multiples. The subsample consists of 29 CTTS (4 of which have upper limits) and 33 WTTS (with no upper limits). We find a probability of $P = 0.05\%-0.07\%$ that the distributions arise from the same parent population, and $\langle \log(L_X/L_\odot) \rangle_{\text{CTTS}} = -3.83 \pm 0.06$ and $\langle \log(L_X/L_\odot) \rangle_{\text{WTTS}} = -3.40 \pm 0.08$. These results are consistent with the results found in the full sample. We can therefore conclude that multiplicity does not influence our
4.4. Results

Figure 4.7: Cumulative distribution of $\log(L_X/L_\odot)$ for CTTS (solid) and for WTTS (dotted).

Figure 4.8: Cumulative distributions of $\log(L_X/L_\odot)$ for WTTS (dotted) and CTTS (solid) for masses smaller than $0.3\, M_\odot$ (left), masses between $0.3$ and $0.7\, M_\odot$ (middle), and masses larger than $0.7\, M_\odot$ (right).

results.

In Fig. 4.8 we plot the Kaplan-Meier estimator for the distribution of $\log(L_X/L_\odot)$ in three different mass ranges: for stars with masses smaller than $0.3\, M_\odot$, between $0.3$ and $0.7\, M_\odot$, and larger than $0.7\, M_\odot$. For the latter two mass bins, the distributions belong to two different parent populations at the $> 94\%$ level. For lower masses ($M < 0.3\, M_\odot$), we find the CTTS and WTTS distributions to be similar, but we still find larger $\log(L_X/L_\odot)$ for WTTS than for CTTS. The difference between the two populations is not significant at the
Figure 4.9: X-ray Luminosity Function (XLF) for CTTS (solid) and WTTS (dotted).

$\approx 10\%$ level possibly because of the small size of the stellar sample in this mass range.

Our results can be compared with the distributions found in the COUP survey, shown in Fig. 16 of Preibisch et al. (2005). In the latter figure, the stars are classified according to the $8542$ Å Ca II line, which is an indicator of disk accretion, similar to the EW(Hα) used in our work. In Orion, a substantial difference has been found between the distributions of accreting and non-accreting stars in the mass ranges $0.2$–$0.3\ M_\odot$ and $0.3$–$0.5\ M_\odot$. However, for $0.5$–$1\ M_\odot$, the two distributions appeared to be compatible, in contrast to our findings that show fainter CTTS consistently in all mass ranges.

### 4.4.2 Correlations between X-Ray Parameters

**The X-ray Luminosity Function**

In Fig. 4.9 we display the X-ray luminosity function (XLF) for WTTS and CTTS for our Taurus sample. The XLF has again been calculated using the Kaplan-Meier estimator in ASURV, so that the few upper limits have also been considered. The total number of sources used was 105, 56 of them being CTTS (including 6 upper limits) and 49 WTTS (including 1 upper limit). The WTTS are again more luminous than CTTS by a factor of about 2 (with mean values $<\log L_X>_C = 29.51$ and $<\log L_X>_W = 29.80$). The probability that the two distributions arise from the same parent population is $7\%$–$10\%$, computed
4.4. Results

Figure 4.10: X-ray Luminosity Function (XLF) for CTTS (solid) and WTTS (dotted) for masses smaller than 0.3 \( M_\odot \) (left), masses between 0.3 and 0.7 \( M_\odot \) (middle), and masses larger than 0.7 \( M_\odot \) (right).

If we restrict the stellar sample to stars with no recognized multiplicity, we obtain average X-ray luminosities of \( \langle \log L_X \rangle_C = 29.38 \) and \( \langle \log L_X \rangle_W = 29.65 \), for samples consisting of 32 CTTS (5 upper limits) and 36 WTTS (1 upper limit). The difference between the two stellar samples is 0.3 dex (i.e., a factor of two), similar to what we found for the full sample. However, the two-sample tests give a larger probability (P = 12\%-33\%) that the two stellar groups arise from the same parent population. Among the multiple sources, 24 are CTTS (with 1 upper limit), but only 13 are WTTS (with no upper limit). By adding the multiples to the sample of single stars, we expect that the distributions slightly shift toward larger \( L_X \), and because there are significantly more multiple CTTS, the CTTS distribution of the total sample should be more similar to the WTTS total distribution, but the opposite trend is seen. We conclude that the trends are grossly the same for the total sample and the single-star subsamples, the larger probability being due to significantly smaller samples that are compared. Overall, thus, CTTS are recognized as being X-ray deficient when compared to WTTS.

Fig. 4.10 shows the X-ray luminosity function for the same three mass ranges as used in Sect. 4.4.1 (\( M < 0.3 M_\odot \) in the left panel, \( 0.3M_\odot < M < 0.7M_\odot \) in the middle panel, and \( M > 0.7M_\odot \) in the right panel). Again, we find the largest difference between CTTS and WTTS for the two higher-mass bins, with probabilities of only 2\%-4\% that the distributions belong to the same parent population. The probability is substantially larger for \( M < 0.3M_\odot \) (29\%-32\%), but the statistics are also considerably poorer.

Considering the difference in the XLFs of CTTS and WTTS alone, a possible cause could be that the bolometric luminosity function of CTTS would indicate lower luminosities \( L_* \) than for WTTS, which would result in lower
average $L_X$ provided that $L_X/L_\star \approx$ constant, i.e., that saturation applies for all stars. In Fig. 4.11, we plot the distributions of $L_\star$ for WTTS and CTTS. In fact, the CTTS are found to be slightly more luminous than the WTTS. We find $\langle \log L_\star \rangle_C = 33.35 \pm 0.08$ and $\langle \log L_\star \rangle_W = 33.19 \pm 0.09$. The probability that the distributions arise from the same parent population is 15–21%, i.e. making the difference marginal. We conclude that because $L_X$ is linearly correlated with $L_\star$ (Fig. 4.5), the difference in $L_X$ for the two samples is intrinsic, which is of course a reconfirmation of our previous finding that the distributions of $L_X/L_\star$ also indicate lower activity for CTTS compared to WTTS.

Absorption

In Fig. 4.12 we plot the distribution of $N_H$ for accreting (solid line) and non-accreting stars (dotted line) calculated using the Kaplan-Meier estimator in ASURV. The logarithmic average of $N_H$ for CTTS is $\langle N_H \rangle_C = 4.2 \times 10^{21}$ cm$^{-2}$ and is more than a factor of two larger than the average for WTTS ($\langle N_H \rangle_W = 1.8 \times 10^{21}$ cm$^{-2}$).

The $N_H$ values found from our spectral fits are roughly consistent with the visual extinctions $A_V$ and the infrared extinctions $A_J$ if we assume a standard gas-to-dust ratio ($N_H/A_V = 2 \times 10^{21}$ cm$^{-2}$ mag$^{-1}$; $N_H/A_J = 7.1 \times 10^{21}$ cm$^{-2}$ mag$^{-1}$; Vuong et al. 2003). For a detailed discussion on the gas-to-dust ratio in TMC, we refer the reader to Glauser et al. (2007, in preparation). High photoelectric absorption might influence the spectral fits to
low-resolution spectra, because the coolest plasma components, more affected by absorption, cannot be reliably quantified.

We studied possible biases introduced by high absorption by correlating $N_H$ with $T_{av}$ and $L_X$. We have found $L_X$ to range between approximately $10^{28}$ erg s$^{-1}$ and $10^{31}$ erg s$^{-1}$, independent of the photoelectric absorption. The uncertainty of the determination of $L_X$, on the other hand, does increase with increasing $N_H$ (and decreasing number of counts in the spectrum) as derived by Güdel et al. (2007a). Similarly, if we correlate $T_{av}$ with $N_H$ (Fig. 4.13), we find a larger range of $T_{av}$ (symmetrically around log $T_{av}$[K]≈ 7.0) for highly absorbed sources, while $T_{av}$ is found to be similar for all the sources with low absorption ($N_H < 10^{21}$ cm$^{-2}$), log $T_{av}$[K]≈ 7.0 ± 0.2. The larger range of $T_{av}$ at higher $N_H$ is likely to be the result of larger scatter due to less reliable spectral fitting. In any case, there is no trend toward higher $T_{av}$ for higher $N_H$.

**Correlation of $L_X$ with Electron Temperature**

In Fig. 4.14 we plot $T_{av}$ as a function of $L_X$ and as a function of the X-ray surface flux ($F_X$) for CTTS and WTTS, respectively. The surface fluxes have been calculated using the radii reported in Table 10 of Güdel et al. (2007a). In the plots for the WTTS we also show values for six main-sequence G-type solar analog stars (Chapter 2), for 5 K-type stars (AB Dor from Sanz-Forcada et al. 2003, and ε Eri, 70 Oph A&B, 36 Oph A&B from Wood & Linsky 2006), and for 6 M-type main-sequence stars (EQ Peg, AT Mic, AD Leo and EV Lac...
Figure 4.13: $T_{av}$ as a function of $N_H$ for our sample.


Spectral fits to low-resolution spectra that are subject to photoelectric absorption tend to ignore the coolest plasma components, as the soft part of the spectrum is most severely affected by the absorption. CTTS are on average more absorbed than WTTS. Given the larger absorptions, there could be a bias toward higher average temperatures in CTTS, although such a trend is not visible in Fig. 4.13. We nevertheless counteract a possible residual bias by restricting the stellar sample used for the correlation to stars with $N_H < 3 \times 10^{21}$ cm$^{-2}$. The logarithmic means of $N_H$ for CTTS and WTTS after these restrictions are $1.2 \times 10^{21}$ cm$^{-2}$ and $8 \times 10^{20}$ cm$^{-2}$, respectively, making these samples very similar with regard to absorption properties.

Further, we exclude the very faint sources (with less than 100 counts collectively in the three detectors) that could also produce unreliable $L_X$ and $T_{av}$ results. In Fig. 4.14 the filled circles represent the stars used for the linear regression fit, while the stars excluded from the fit are plotted with small crosses. Overall, the absorbed and faint sources fit well to the trends found from less absorbed and more luminous sources, but their scatter tends to be larger.

For CTTS, we find almost no correlation between $T_{av}$ and $L_X$ or $F_X$ (the correlation coefficients are 0.06 and 0.11, respectively). On the contrary, for WTTS $T_{av}$ is clearly correlated with both $L_X$ and $F_X$. The correlation coefficients are 0.69 and 0.72 for $L_X$ and $F_X$, with 33 and 32 data points, respectively (Table 4.1). The probability that no correlation is present is $< 1\%$ in either case. We computed the linear regression using the bi-
4.4. Results

Figure 4.14: Left panel: $T_{\text{av}}$ as a function of $L_X$ for CTTS (top) and WTTS (bottom). Right panel: $T_{\text{av}}$ as a function of $F_X$ for CTTS (top) and WTTS (bottom). The low-absorption TTS samples are marked by filled black bullets, while small crosses give loci of high-absorption objects or sources with few counts (see text for details). Black diamonds mark solar analog stars (Chapter 2) and blue and green diamonds mark K- and M-type main-sequence stars, respectively (see text for references). The straight lines in the WTTS plots are linear regression fits (based on bisector regression, the dashed lines illustrating the error ranges in the slopes).

sector OLS algorithm (no a priori relation between the two measured variables assumed) to find $\log T_{\text{av}} = (0.23 \pm 0.03) \log L_X + (0.13 \pm 0.93)$ and $\log T_{\text{av}} = (0.26 \pm 0.03) \log F_X + (5.21 \pm 0.24)$. WTTS follow a trend similar to that shown by MS stars in the $T_{\text{av}}$ vs. $L_X$ relation. In the $T_{\text{av}}$ vs. $F_X$ relation, on the other hand, we find that WTTS are in general hotter than MS stars for a given $F_X$.

We have checked these results using the EM algorithm, finding slightly shallower slopes (see Table 4.1). Shallower slopes are expected in the EM algorithm when compared to the bisector OLS algorithm (Isobe et al. 1990). For CTTS, where no correlation is found, the two algorithms result in completely different
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Figure 4.15: Kaplan-Meier estimator for $T_{av}$ for CTTS (solid) and WTTS (dotted). Only sources with more than 100 counts in the three EPIC detectors are used.

slopes (Table 4.1), an indication of absence of a linear regression (Isobe et al. 1990; the different slopes in the absence of a correlation are a consequence of the defining minimization of the algorithm. The EM algorithm returns a slope of $\approx 0$, whereas the bisector algorithm yields a slope around unity).

We use the Kaplan-Meier estimator in ASURV to compare the distributions of $T_{av}$ for CTTS and WTTS. The two distributions are shown in Fig. 4.15 for the case where we do not apply a restriction to $N_H$: the solid line represents CTTS, while the dotted line represents WTTS. Only stars with more than 100 counts in the three EPIC detectors are used. The probability that the two distributions arise from the same parent population is 0.7–2%. If we restrict the sample to stars with low $N_H (< 3 \times 10^{21} \text{ cm}^{-2})$, we find the mean value of $\log T_{av}[K] = 7.10$ with $\sigma = 0.22$ for CTTS and $\log T_{av}[K] = 6.88$ with $\sigma = 0.17$ for WTTS. The distribution is similar to the one found in Fig. 4.15 and the probability for the CTTS and WTTS distributions to originate from the same parent population is again only 0.7–2%.

We have checked if the difference found in the plasma temperatures of WTTS and CTTS could be attributed to abundance anomalies that may not have correctly been accounted for in the spectral fits. Kastner et al. (2002), Stelzer & Schmitt (2004), and Drake et al. (2005b) found large Ne/Fe and Ne/O abundance ratios in the spectrum of TW Hya. We have therefore fitted the spectra of the 19 CTTS with $N_H < 3 \times 10^{21} \text{ cm}^{-2}$ and more than 100 counts in the combined EPIC spectra (filled black bullets in Fig. 4.14)
adopting abundances as found in TW Hya (O = 0.2, Ne = 2.0, Fe = 0.2, with respect to the solar photospheric abundances of Anders & Grevesse 1989, all other abundances as given in Sect. 4.3). The average temperatures obtained with this model are generally consistent within 0.1 dex with the temperatures found based on our standard abundances. Only for one star, HO Tau AB, did we find an average temperature significantly lower, while the general trend toward higher $T_{\text{av}}$ for CTTS remains unchanged. We can therefore exclude that the difference in temperatures is induced by abundance anomalies as those observed in TW Hya.

In Chapter 3 we derived the thermal structure of nine pre-main sequence stars from XEST based on high-resolution Reflection Grating Spectrometer data, using variable abundances. They found $T_{\text{av}}$ to be compatible with values used here, which were derived from EPIC CCD spectra (an exception is the CTTS SU Aur, for which the temperature found with RGS is even higher than that derived from the EPIC spectra). In Chapter 3, however, a difference in the abundances has been found between stars of spectral type K and stars of spectral type G. The abundances found for the K-type stars reflect approximately the abundances used for the XEST EPIC fits (following an inverse FIP effect), while G-type stars show lower Ne/Fe and O/Fe abundance ratios. We therefore fitted the four G-type stars in our stellar sample with an abundance pattern as found for this spectral class in Chapter 3. Again, we did not find a significant change in temperatures. In summary, we do not find any appreciable effect that abundance anomalies other than those adopted in our study might have on the temperature determination. We also note that Scelsi et al. (2006, in preparation) studied the abundances derived from the EPIC spectra of the brightest sources in the XEST sample and found average abundances very similar to the standard abundances used in our CCD fits.

The difference in the coronal temperatures of CTTS and WTTS is in particular due to the larger $T_{\text{av}}$ found in the CTTS with low $L_X$. We therefore calculated the mean of all $T_{\text{av}}$ values for CTTS and WTTS with $L_X < 3 \times 10^{29}$ erg s$^{-1}$, low absorption ($N_H < 3 \times 10^{21}$ cm$^{-2}$), and more than 100 counts in the three detectors. For these stars, we computed the errors in $T_{\text{av}}$ as follows: We determined the 68% confidence contour on the $\beta - T_0$ plane for these two parameters of interest, i.e. the loci for which a fit can be achieved whose $\chi^2$ is larger by $\Delta \chi^2 = 2.3 \ (1 \sigma)$ than the $\chi^2$ of the best fit. We then found the minimum and the maximum $T_{\text{av}}$ for this subset of models, and thus defined the error range for $T_{\text{av}}$. Using these errors, we computed the weighted mean of $\log T_{\text{av}}$. For CTTS, we neglected DD Tau AB, which shows extraordinarily high temperatures in two different observations (two bullets at the hottest temperature in Fig. 4.14). For CTTS, we find $\langle \log T_{\text{av}} \rangle = 6.97 \pm 0.06 \ (6.98 \pm 0.06 \text{ if DD Tau AB is also considered})$ while for WTTS we find $\langle \log T_{\text{av}} \rangle = 6.81 \pm 0.05$. We therefore conclude that the CTTS and WTTS with $L_X < 3 \times 10^{29}$ erg s$^{-1}$
are different at a 3σ level, fully supporting the significant differences in the regression fits that are based on the entire \( L_X \) range.

In conclusion, we find the CTTS X-ray sources to be hotter than WTTS at a confidence level of \( \gtrsim 98\% \), and this result is partly due to the presence of a \( L_X - T_{av} \) relation for WTTS but its absence in CTTS. Further, the WTTS relation coincides with relations valid for main-sequence stars of different spectral types, including saturated and non-saturated stars at different evolutionary stages.

4.5 Discussion

4.5.1 Summary of Trends

We now discuss the trends and correlations described in the previous section and will also put them into a context with previous reports, in particular from the COUP project.

The most significant correlations that we reported above are those between stellar mass and \( L_X \) (slope \( \approx 1.7 \)), between stellar bolometric luminosity \( L_s \) and \( L_X \) (slope \( \approx 1 \)), and between \( L_X \) and average electron temperature \( T_{av} \) (slope 0.15–0.23), the latter applying only to WTTS.

Further, we have found that \( L_X \) and \( L_X/L_s \) are both lower, on average, for CTTS than for WTTS, each by a factor of \( \approx 2 \), compatible with the finding that the distributions of \( L_s \) are similar for the two samples. In contrast, \( T_{av} \) is, on average, higher by a factor of \( \approx 1.7 \) for CTTS than for WTTS.

Finally, we have studied possible correlations between \( L_X, T_{av}, \) or \( L_X/L_s \) and the accretion rate but found at best unconvincing correlations. The same is true for a trend between \( L_X \) and age.

4.5.2 Comparison with Previous Studies

The \( L_X - \text{mass} \) correlation. This relation has been reported prior to \textit{XMM-Newton} and \textit{Chandra} studies of star-forming regions, but with largely varying regressions. Feigelson at al. (1993) found a slope of 3.6 ± 0.6 for a sample of low-mass stars in the Chamaeleon I dark cloud based on ROSAT observations. There may be problems with more numerous upper limits at the low-mass (and low-luminosity) end of the distribution in this study, as noticed by Preibisch et al. (2005). On the other hand, the COUP sample reveals a very similar correlation to ours, with a slope only marginally smaller (1.44 ± 0.10) than for XEST (1.69 ± 0.11). The TMC sample thus essentially confirms the COUP results, and the residual difference might be due to a somewhat different distribution of stars in the HRD, perhaps indicating a different age distribution, as suggested from our discussion of this relation below.
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A clear difference between the two studies is seen in the scatter around the regression curves. While Preibisch et al. (2005) report a standard deviation of 0.65 dex (factor of 4.5) around the best-fit line, we find for the XEST sample values between 0.38 dex (factor of 2.4, for WTTS) and 0.45 dex (factor of 2.8, for CTTS). We do not have a clear explanation for the smaller scatter in XEST, but note that i) similar findings apply to other correlations discussed below, and ii) the scatter found in the XEST results is close to the intrinsic uncertainties of any $L_X$ measurement of magnetically active stars as these commonly vary by such factors on various time scales; Preibisch et al. (2005) give a factor of 2 variation on long (yearly) time scales for the Orion sample. We are thus confident that the quality of our mass – $L_X$ correlation corresponds to the minimum scatter that must be expected from snapshot observations of magnetically active stars.

The $L_X - L_*$ correlation. The linear correlation between these variables expresses the classical result of X-ray saturation that has been found empirically for main-sequence stars (Vilhu & Rucinski 1983). A similar law applies to very active main-sequence and subgiant stars (see review by Güdel 2004a and references therein), and certainly also to pre-main sequence stars at various evolutionary stages (Flaccomio et al. 2003b). Again, the COUP study is in complete agreement with our results, its regression slope being 1.04 ± 0.06. There is, however, a significant difference between the XEST and the COUP studies once CTTS and WTTS are treated separately. Preibisch et al. (2005) find a well-defined linear correlation for WTTS (standard deviation around best-fit regression of 0.52 dex), while for CTTS the scatter dominates (standard deviation = 0.72 dex) and the relation is significantly flatter. The CTTS data points span a range of 3 orders of magnitude at a given $L_*$. In XEST, the standard deviation of the scattered points is only ≈ 0.4 dex for CTTS, WTTS, and the entire sample, with a range of $L_X$ values at a given $L_*$ of about 1.5 dex. We are again not in a situation to explain the much tighter correlations for the XEST survey, but note that our spectral-fit methodology may suppress numerical uncertainty introduced by photoelectric absorption that suppresses evidence of plasma components at lower temperatures. Because we used an emission-measure distribution with a prescribed low-temperature shape as usually found in magnetically active stars, the presence of the coolest components is interpreted based on the presence of well-detected hotter plasma. An error analysis for $L_X$ based in particular on $N_H$ shows that for 76% of the sources in XEST, the intrinsic error range in $L_X$ due to $N_H$ is smaller than a factor of 3 (0.5 dex, in fact mostly much smaller), and the largest errors are obtained for faint sources subject to $N_H > (2 - 3) \times 10^{21}$ cm$^{-3}$ (Güdel et al. 2007a). We note that Preibisch et al. (2005) used an X-ray luminosity averaged over the 10 days of exposure by Chandra, while in the XEST sample we have neglected time intervals containing obvious flares. However, Preibisch et al. (2005) found that the average $L_X$ and the quiescent (“characteristic”) $L_X$ differ by a median
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factor of 0.78, which would not be sufficient to explain the large scatter found in their $L_X - L_*$ correlation for CTTS.

Comparison with main-sequence stars. Preibisch et al. (2005) also compare their $L_X - L_*$ correlation with field main-sequence stars, and find a much shallower slope for the latter, but also a very large scatter. They similarly compare the $L_X - L_*$ mass relations with field stars, but there, they find a similar slope in the regression. We will not perform this comparison here, for the following reason. Field stars are found at various evolutionary stages, and as a consequence of stellar spin-down with age, the X-ray activity is subject to an evolutionary decay. Solar analogs decrease in $L_X$ by three orders of magnitude from the zero-age main sequence to the end of the main-sequence life (Güdel et al. 1997a; Chapter 2). Further, the evolutionary speed is different for G stars and low-mass M dwarfs, the latter remaining at relatively high activity levels for a longer time (see Figs. 40 and 41 in Güdel 2004a). Much of the scatter in $L_X$ for a given mass or a given $L_*$ is thus due to mass-dependent evolutionary decay, and any trend in $L_X$ vs. $L_*$ depends on the stellar age distribution. In contrast, both in active main-sequence stars and TTS, no evolutionary effects are expected for the $L_X - L_*$ relation if the stars are in a saturated regime, and therefore $L_X \propto L_*$. For the $L_X - L_*$ mass relation, the scatter in $L_X$ for a given mass is only about one order of magnitude; this scatter is indirectly due to the scatter in $L_*$ in the sample, due to different ages of stars of similar mass that contract vertically along the Hayashi track, provided that the X-ray emission remains in a saturated state (see HRD in Fig. 11 in Güdel et al. 2007a). The scatter in $L_X$ due to evolution on the main sequence is much larger (3 orders of magnitude) and is due to intrinsic decay of the dynamo due to stellar spin down when the X-ray emission is no longer in a saturated state.

The $L_X - T_{av}$ correlation. A dependence between coronal electron temperature and emission measure or $L_X$ (or normalized quantities such as the specific emission measure or surface X-ray flux) has first been noted by Vaiiana (1983) and Schrijver et al. (1984). Quantitatively, for solar analogs, $L_X \propto T^{4.5\pm0.3}$ (Güdel et al. 1997a, see Güdel 2004a for a review). For the pre-main sequence sample in the COUP survey, Preibisch et al. (2005) report a steep increase of the X-ray surface flux with the hotter temperature of their 2-component spectra, namely $F_X \propto T_2^5$. On the other hand, they find a relatively constant lower temperature, namely $T_1 \approx 10$ MK. In our study, we apply a more physically appropriate continuous emission measure distribution that does not distinguish between two isothermal components but that shows two power-law slopes on either side of the peak. The distribution of the logarithmically averaged temperatures (Fig. 4.15) does not show a preferred value but a smooth distribution in the range 4–30 MK around a mean of 7.6 MK for WTTS and 12.6 MK for CTTS (Sect. 4.4.2). Our regression curve for the $L_X - T_{av}$ indicates $L_X \propto T_{av}^{4.3\pm6.7}$ and $F_X \propto T_{av}^{3.8\pm5.6}$, compatible with the solar-
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analog relation as well as with the COUP relation for $T_2$. We note, however, that we find this trend only for WTTS (no separate analysis was provided for COUP).

$log(L_X/L_\ast)$ distributions. Our distributions show the fractional X-ray luminosity of CTTS to be suppressed by a factor of $\approx 2$ compared to WTTS, with a mean $log(L_X/L_\ast) \approx -3.39 \pm 0.06$ and $-3.73 \pm 0.05$ for WTTS and CTTS, respectively. These values agree excellently with those of COUP: $log(L_X/L_\ast) \approx -3.31$ and $-3.73$ for WTTS and CTTS, respectively (Preibisch et al. 2005). XEST contrasts with COUP in that CTTS are less X-ray efficient for all considered mass bins, whereas Preibisch et al. (2005) found no difference in the ranges 0.1-0.2 $M_\odot$ and 0.5-1 $M_\odot$.

4.5.3 The $L_X - \text{Mass Relation}$

Among the clearest correlations we have identified is the $L_X$-mass relation that closely corresponds to the finding in the COUP study. We now test the following: If we assume that TTS are in a saturated state (i.e. $L_X \propto L_\ast$, Fig. 4.5) and a relation between $L_\ast$ and stellar mass exists (related to the age distribution of the stars and details of the evolution along the pre-main-sequence tracks), then the relation between $L_X$ and mass could simply be a consequence of these two relations. Main-sequence stars follow the well-known mass-bolometric luminosity relation, which for stars in the mass range of 0.1–1.5 $M_\odot$ reads $L_\ast \propto M^{3.0}$ (from a regression analysis using the Siess et al. 2000 ZAMS data). For pre-main sequence stars, a mass-bolometric luminosity relation is not obvious; during the contraction phase, a star of a given mass decreases its $L_\ast$ by up to 2 orders of magnitude. However, if most stars in a sample show similar ages, then an approximate mass-$L_\ast$ relation may apply to the respective isochrone. Fig. 4.16 illustrates the measured relation between $L_\ast$ and mass. The relation is rather tight, with a correlation coefficient of 0.90 for 113 sources. The Y/X OLS regression gives $log L_\ast/L_\odot = (1.49 \pm 0.07) log M + (0.23 \pm 0.04)$, with a standard deviation of 0.31. This relation can be compared with the theoretical prediction for an average isochrone appropriate for the XEST sample. We found that the logarithmically averaged age of our targets is 2.4 Myr. Adopting the Siess et al. (2000) isochrones, we find, from a linear regression fit, a dependence $L_\ast \propto M^{1.24}$, i.e., similar to the observed dependence and thus supporting our interpretation. We note that the XEST sample is of course not located on an isochrone (see Fig. 11 in Gudel et al. 2007a), and that other evolutionary calculations may lead to somewhat different slopes of the isochrones.

Adopting $log(L_X/L_\ast) = -3.5$ for our entire TTS sample (see Fig. 4.5), we infer a relation $log L_X = 1.49 log M + 30.31$, similar to the correlation found in Sect. 4.4.1. In Fig. 4.17 we plot $L_X$ as a function of mass after normalizing the observed $L_X$ with $L_X$ predicted from the above formula. The correlation
Figure 4.16: Stellar bolometric luminosity as a function of mass. The straight line is a Y/X OLS regression fit; dashed lines illustrate the 1-sigma errors in the slope.

Figure 4.17: $L_X$ as a function of mass after renormalizing $L_X$ with the expected $L_X$ based on mass-bolometric luminosity relation and the saturation law (see text for details). Different symbols mark different types of stars.
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found in Fig. 4.1 now disappears completely. The scatter in Fig. 4.17 is due to the scatter in Fig. 4.5 and in Fig. 4.16, i.e., intrinsic scatter not due to \( L_\star \) or \( M_\star \), but for example due to the evolutionary decrease of \( L_\star \) and \( L_X \) along the Hayashi track for a given mass.

We therefore conclude that the \( L_X \)-mass relation is not an intrinsic relation but a consequence of an approximate mass-luminosity relation for stars with similar ages, combined with a saturation law.

Comparing the stars in the HRD of the TMC (Fig. 11 in Güdel et al. 2007a) with the HRD of the ONC sample (Fig. 1 in Preibisch et al. 2005), we note that the ONC sample is somewhat younger, although it is not tightly arranged along one isochrone, but tends to show a somewhat flatter slope than the TMC sample. The slightly shallower \( L_X \)-mass relation in Preibisch et al. (2005) may therefore be a straightforward consequence of the younger average age of the ONC sample.

4.5.4 Origin of the X-Ray Emission: CTTS vs WTTS

The question that we address in this section is on the origin of the X-ray emission. How do CTTS and WTTS differ, and what may be the causes? Where and how are the X-rays formed?

The relevant relations we have identified in this chapter are the following: i) CTTS show, on average, a smaller X-ray luminosity in the EPIC band; ii) CTTS also reveal a significantly lower fractional X-ray luminosity, \( L_X/L_\star \), than WTTS. iii) The average electron temperature in WTTS correlates with the total \( L_X \) for WTTS, but this is not the case for CTTS; the CTTS temperatures are on average significantly higher than those of WTTS.

Evidence for Coronal Emission

The bulk of the X-ray emission described in this chapter and in the COUP survey is consistent with an origin in a magnetic corona. The overall temperatures measured in TTS and the X-ray luminosities are similar to values also found in extremely active main-sequence and subgiant stars (e.g., review by Güdel 2004a). Also, frequent flaring in TTS (Wolk et al. 2005; Stelzer et al. 2007; Franciosini et al. 2007) clearly points to a coronal (or magnetospheric) origin of the X-ray emission. Although many of the T Tauri stars in the XEST sample are thought to be fully convective and hence unable to support a solar-like \( \alpha \)\( \Omega \)-dynamo, fully-convective main-sequence stars do show magnetic activity, and dynamo mechanisms have been proposed that may operate in such stars (e.g., Dobler et al. 2006 and references therein; Küker & Rüdiger 1999). A discussion of what XEST can tell us about the dynamos acting in T Tauri stars in Taurus-Auriga is given by Briggs et al. (2007).
We have found in the previous section that the $L_X$-mass relation is due to saturation and a mass-bolometric luminosity relation for TMC stars. The most fundamental relation is indeed the one between $L_X$ and $L_*$, which indicates that the PMS stars in Taurus are saturated at $\log(L_X/L_*) \approx -3.5$, in analogy to main-sequence stars. This result suggests that the bulk of X-ray emission in PMS stars by CCD detectors used in XEST and in COUP arises from a magnetic corona, and we therefore suggest that similar processes as in MS stars should be responsible for the dominant energy output from the hot plasma (we will rediscuss emission from the softest X-ray emitting plasmas below).

Is a Lower $L_X$ Intrinsic to CTTS Coronae?

We have tested (Sect. 4.4.2) whether the lower $L_X$ seen in CTTS compared to WTTS could be a consequence of generally lower $L_*$ of CTTS, provided that all stars are in a similar saturation regime. This can be explicitly rejected because we found the $L_*$ distributions for the two samples to be similar, with a trend that it is rather the CTTS sample that is slightly more luminous. Also, the $L_X/L_*$ distributions are explicitly different, offset by about the same factor as the $L_X$ distributions themselves.

In previous works the lower X-ray activity of CTTS compared to WTTS in Taurus-Auriga has been attributed to the slower rotation of CTTS and an anticorrelation of activity and rotation period as exhibited by active solar-like stars (e.g., Neuhäuser et al. 1995). However, the lower X-ray activity of CTTS has been observed in other star-forming regions where an anticorrelation of activity and rotation period is clearly not seen (e.g., Preibisch et al. 2005). Briggs et al. (2007) demonstrate that an apparent activity-rotation relation in Taurus-Auriga naturally results from the dependences of activity on mass and accretion status reported here and in other star-forming regions because the fast rotators in Taurus-Auriga are mainly higher-mass and non-accreting while the slow rotators are mainly lower-mass and accreting. There is no convincing evidence for an anticorrelation of X-ray activity and rotation period in T Tauri stars, and therefore no evidence that the lower activity of CTTS is due to their slower rotation.

Indeed, even if a solar-like dynamo operates in T Tauri stars, their long convective turnover timescales lead to the expectation that all stars with measured rotation periods should have saturated (or supersaturated) emission (e.g., Preibisch et al. 2005) and show no anticorrelation of activity and rotation period.

Different internal structure in WTTS and CTTS (unless induced by accretion processes) is not a likely explanation either. As was recognized in early infrared and optical surveys of the TMC stellar population, WTTS and CTTS occupy the same region in the HRD (Kenyon & Hartmann 1995), with no evolutionary separation, indicating that the transition from CTTS to WTTS
occurs at very different ages for different stars. This suggests that the lower $L_X$ of CTTS is an intrinsic property of a corona that is heated in the presence of an accretion disk and active accretion onto the star. We now briefly consider the influence of accretion on coronal heating.

**The Role of Active Accretion**

Accretion has variously been suggested to enhance or to suppress plasma heating. First, accretion hot spots may heat plasma to temperatures in excess of one million K as gas shocks near the surface in nearly free fall (e.g., Calvet & Gullbring 1998). High-resolution X-ray spectroscopy has provided some indirect evidence that such accretion-induced X-rays might constitute an important part of the measurable spectra. Kastner et al. (2002) have interpreted exceptionally high-densities and very cool ($T \approx 3$ MK) X-ray emitting material in the CTTS TW Hya as being the result of accretion shocks. This idea was further elaborated by Stelzer & Schmitt (2004) who also suggested that anomalously high Ne and N abundance in this X-ray source indicates that refractory elements such as Fe condense onto dust grains in the disks, and that this material is eventually not accreted. A small number of additional X-ray spectra have been studied, but the situation is complex and contradictory: Schmitt et al. (2005) and Robrade & Schmitt (2006) report intermediately high densities in BP Tau ($n_e \approx 3 \times 10^{11}$ cm$^{-3}$), and Günther et al. (2006) find similarly high densities in V4046 Sgr. On the other hand, very low densities and no strong abundance anomalies have been found in T Tau (Güdel et al. 2007b) and the accreting Herbig star AB Aur (see Chapter 4). However, in Chapter 3 we found evidence that CTTS in general maintain an excess of cool (1-4 MK) plasma compared to WTTS. This expresses itself in O$^7$ vii line fluxes that are similar to the flux in the O$^8$ viii Ly$\alpha$ line, a condition very different from WTTS where the O$^7$ vii lines are very faint (similar to magnetically active ZAMS stars).

This soft excess seems - as far as the still small statistical sample of stars suggests - to be related to accretion in CTTS. But where are the X-rays produced? For T Tau and AB Aur, a production in accretion shocks is unlikely (Güdel et al. 2007c, Chapter 3) while accretion shocks may be responsible for the softest emission in TW Hya, BP Tau, and V4046 Sgr (Kastner et al. 2002; Stelzer & Schmitt 2004; Schmitt et al. 2005; Günther et al. 2006).

**Coronal Modification by Accretion Streams?**

An alternative possibility is that accretion influences the magnetic field structure of the stars, or that the accreting material is changing the heating behavior in the coronal magnetic fields. Preibisch et al. (2005) proposed that the mass-loaded loops of accreting stars are denser than the loops of WTTS, so that
when a magnetic reconnection event occurs, the plasma would be heated to much lower temperatures outside the X-ray detection limit. This could account for the deficiency of X-ray luminosity in CTTS compared to WTTS. A similar scenario has been proposed by Güdel et al. (2007b) to specifically explain the soft excess observed in T Tau. In this scenario, the magnetospheric geometry is influenced by the accretion stream. A fraction of the cool accreting material enters the coronal active regions and cools the magnetic loops there for three reasons: i) the accretion flow may reorganize magnetic fields, stretching them out and making them less susceptible to magnetic reconnection; ii) the accreting material is cold, lowering the resultant temperature on the loops when mixing with the hot plasma; and iii) the accreting material itself adds to the coronal density, inducing larger radiative losses and more rapid cooling. These resulting soft X-rays from cool material can be detected by the RGS instruments that are sensitive at low energies and provide the spectral resolution to record flux from lines formed at temperatures below 3 MK, such as O\textsuperscript{vii} and N\textsuperscript{vi}, but the same emission will not be separately identified by the EPIC detectors that are relatively insensitive at the relevant temperature, and provide only very low energy resolution. Güdel et al. (2007b) estimated for T Tau that only of order 1% of the accreting material would be needed to penetrate active regions on the star and be heated to 2 MK.

If the total coronal energy release rate (averaged over time scales longer than energy release events such as flares) is determined by the stellar dynamo that forms a magnetic corona (and by convective properties near the stellar surface), then we would expect similar coronal radiative losses for CTTS and WTTS. Could it be that the radiative output from the corona is indeed equivalent but that part of the coronal emission has shifted to the softest part of the spectrum, remaining undetected in EPIC CCD spectroscopy while detected as a soft excess by RGS? At least in the case of T Tau, the 0.3-10 keV X-ray luminosity has been severely underestimated by EPIC CCD analysis alone, as these instruments missed the softest component, also subject to considerable photoelectric absorption, and suggested $L_X$ to be only $\sim 60\%$ of $L_X$ determined from the combined RGS+EPIC spectra (Güdel et al. 2007b).

We systematically studied our RGS spectra (of Chapter 3) to find out whether the soft excess in our CTTS sample provides the “missing luminosity”. We have therefore compared the unabsorbed 0.1-0.5 keV X-ray luminosity from the EPIC spectral fits (Güdel et al. 2007a) with $L_X$ in the same range from the combined EPIC+RGS fits (Chapter 3). The comparison is useful only for targets that do not suffer from strong absorption; this is the case for the CTTS BP Tau and DN Tau and the WTTS HD 283572, V773 Tau, and V410 Tau. We found no systematic difference to explain the factor of 2 underluminosity of CTTS. It appears that the EPICs record the soft emission from the coolest plasma sufficiently well to register similar $L_X$ as the RGS detectors, but the temperature discrimination is clearly inferior to high-resolution spectroscopy.
Also, the softest range is still dominated by continuum emission from hotter plasma, and the soft excess in these stars provides relatively little spectral flux.

We have found, on the other hand, that CTTS show, on average, higher electron temperatures (averaged over the components detected by the EPIC cameras) than WTTS. This could be an effect of depletion of the intermediate and cooler temperature ranges by the accretion process, as suggested above, thus moving the average temperature of the detected coronal components to higher temperatures. We have therefore tested whether the harder portion of the EPIC spectra which is radiated by the hottest coronal plasma components also shows a statistical difference between CTTS and WTTS. We chose the 1.5–10 keV range for this test. We found, however, the same discrepancy between the two stellar groups, suggesting that the emission measures of the hottest plasma components themselves are also suppressed in CTTS compared to WTTS.

We extended our comparison to the XEST results obtained from EPIC only, but again found CTTS to be underluminous by similar factors in different X-ray energy bands. To explain the deficit of X-ray emission in CTTS, it could thus be that the accretion process is cooling active-region plasma to an extent that it is also no longer detected in the RGS band.

Coronal Heating in WTTS and CTTS due to Flares?

This brings us to the correlation between average coronal temperature and $L_X$ which is present in WTTS but absent in CTTS. WTTS show a trend in which $T_{av}$ increases with $L_X$, and this trend is the same as previously found for main-sequence solar analogs (Güdel et al. 1997a; Chapter 2). We note that this relation remains valid even for stars with different X-ray saturation limits and different $L_\nu$ (see Fig. 4.14). M dwarfs at the saturation limit reveal lower coronal temperatures than K or G dwarfs at their saturation limit. The cause for this relation is not clear. Güdel (2004a) pointed out that the slope of the regression function (using emission measure instead of $L_X$) is the same as the slope of the regression between peak temperature and peak emission measure in stellar flares. Güdel (2004a) hypothesized that coronal emission is formed by a superposition of continuously occurring “stochastic flares”, with the consequence that larger, hotter flares that occur more frequently in more active stars not only produce the dominant portion of the observed emission measure, but also heat the observed plasma to higher temperatures than in lower-activity stars. The larger rate of large flares would be a consequence of denser packing of magnetic fields in more active stars, inducing more frequent explosive magnetic reconnection, including larger flares than in low-activity stars (Güdel et al. 1997a). A similar trend for WTTS as for main-sequence stars is therefore perhaps not surprising: X-ray production in both types of stars is thought to be entirely based on the magnetic field production by the
internal dynamo. This analogy fully supports solar-like coronal processes in WTTS.

The \( L_X - T_{av} \) correlation is absent in CTTS. The distinguishing property of CTTS is active accretion, which thus is most likely the determining factor for the predominantly hot coronal plasma. Temperatures like those determined as \( T_{av} \) in the XEST survey cannot be produced in accretion shocks, again pointing to a predominantly coronal origin of the hot, dominant plasma component. If the flare-heating concept has merit in CTTS as well, then it seems that flares in CTTS are predominantly hot, even if \( L_X \) is low. We can only speculate about the origin of this feature. A possibility are star-disk magnetic fields, but it is unclear why flares occurring in such loop systems should be hotter.

### 4.6 Summary and Conclusions

We have studied X-ray parameters of a large sample of X-ray spectra of CTTS and WTTS in the Taurus Molecular Cloud. Our principal interest has been in a characterization of X-rays in the two types of stars, in finding correlations between X-ray parameters and fundamental stellar properties and among themselves, and most importantly in comparing our findings between accretors and non-accretors. This study has been motivated by numerous previous reports on correlations and differences between CTTS and WTTS in nearby star-forming regions, and in particular by the COUP study of the Orion Nebula Cluster. The XEST project has provided the deepest and, for the surveyed area, most complete X-ray sample in the Taurus region to date. We have used a CTTS and a WTTS sample of comparable size.

We have correlated \( L_X \) and the average coronal temperature, \( T_{av} \), with various stellar parameters, and conclude the following from our study:

- The X-ray luminosity is well correlated with the stellar mass, with a dependence \( L_X \propto M^{1.7} \), similar to what has been shown in COUP and previous TTS studies, but we find that this correlation is only an expression of saturation and a mass-(bolometric) luminosity relation for our pre-main sequence sample. As long as the stellar sample is saturated, \( L_X \) is a function of \( L_* \), and the latter is correlated with stellar mass for a given isochrone. From stellar evolution calculations (e.g., Siess et al. 2000), the functionality between mass and \( L_* \) can be derived. This is fully analogous to main-sequence stars where approximately \( L_* \propto M^3 \) holds. For a typical isochrone of TMC stars with ages of 2–3 Myr, the exponent is smaller as can be seen on a pre-main sequence HRD (Fig. 11 in Güdel et al. 2007a). For our sample, \( L_* \propto M^{1.5} \).

- A saturation relation holds for both CTTS and WTTS, although \( L_X/L_* \) is, on average, smaller by a factor of 2 for CTTS compared to WTTS.
4.6. Summary and Conclusions

- We find that the distributions of $L_x$ are similar for CTTS and WTTS. As a consequence, we find a significant difference in the X-ray luminosity functions for CTTS and WTTS, the former being fainter by about a factor of two. The suppressed X-ray production in CTTS is thus intrinsic to the source and not due to selection bias.

- We emphasize that the lower X-ray production in CTTS refers to the range of plasma temperatures accessible by CCD cameras such as those used here and in COUP. It is possible that some of the energy release is shifted to lower temperatures outside the range easily accessible to CCD detectors. Those soft regions of the X-ray spectrum are also subject to increased photoelectric absorption, which makes detection of cool plasma more difficult. CTTS are indeed more absorbed than WTTS, namely by a factor of $\approx 2.5$.

- We investigated whether all X-ray spectral ranges show X-ray suppression in CTTS. The hardest portion (1.5–10 keV) shows the X-ray deficiency in CTTS vs. WTTS independently, even though CTTS reveal higher average temperatures. We hypothesized that a fraction of the emission measure has been cooled to poorly detectable or undetectable temperatures in CTTS. CTTS indeed show a soft excess in their high-resolution X-ray spectra, characterized by unusually strong O\text{vii} lines from cool plasma (Chapter 3). These lines cannot be resolved by EPIC. We therefore checked whether RGS spectroscopy of little absorbed CTTS stars in Chapter 3 (DN Tau, BP Tau) indicates a relative increase of $L_X$ in the soft 0.1–0.5 keV band relative to WTTS, but found no significant effect. The soft flux may have been sufficiently well detected by EPIC in these low-absorption stars (but with little temperature discrimination). Also, the softest range is still dominated by continuum emission from hotter plasma, and the soft excess in these stars provides relatively little spectral flux. The X-ray deficiency in CTTS thus remains.

The situation is clearly different in T Tau (Güdel et al. 2007b): in this much more strongly absorbed source, a very large amount of very cool X-ray emitting plasma was detected based exclusively on anomalously strong O\text{vii} line emission in the grating spectrum but went unnoticed in CCD spectroscopy. The analysis of the latter spectrum alone led to an underestimate of the 0.3–10 keV luminosity by 40%.

In conclusion, it seems that the entire X-ray range accessible to CCD spectroscopy reveals suppressed X-ray emission compared to WTTS, although additional components may be present at cool temperatures that may be missed by the CCD spectra, especially if $N_H$ is sufficiently high.

- A possible cause for the suppression of X-ray emission in CTTS may be the accretion streams themselves. If only a small portion of the accreting
matter penetrates into hot coronal magnetic structures, the plasma may cool as more matter needs to be heated and as the increase in density increases the cooling efficiency. This may lead to a soft excess (Güdel et al. 2007b), or to a cooling of plasma to temperatures outside the X-ray regime (Preibisch et al. 2005), so that a significant deficiency of X-ray emission may be measured in the spectral range that is accessible to CCD cameras, and that is not subject to significant photoelectric absorption. The soft excess and the hot-plasma deficiency seem to be related to the presence of accretion.

- X-ray production in shocks at the base of accretion streams has been suggested previously from high-resolution spectra of CTTS. The shocked plasma would add soft emission to the spectra as well, but again, CCD spectroscopy may miss this emission, or the latter may be subject to absorption. Our CCD survey does not provide the appropriate means to test X-ray production in accretion shocks in CTTS, and can therefore also not exclude such mechanisms. High-resolution grating spectroscopy is required.
Chapter 5

The First High-Resolution X-Ray Spectrum of a Herbig Star: AB Aurigae

ABSTRACT: The X-ray emission from Herbig Ae/Be stars remains to be explained. In later-type T Tauri stars, X-rays are thought to be produced by magnetically trapped coronal plasma, although accretion-shock induced X-rays have also been suggested. In earlier-type (OB) stars, shocks in unstable winds are thought to produce X-rays. We present the first high-resolution X-ray spectrum of a prototypical Herbig star (AB Aurigae), measure and interpret various spectral features, and compare our results with model predictions. We use X-ray spectroscopy data from the XMM-Newton Reflection Grating Spectrometers and the EPIC instruments. The spectra are interpreted using thermal, optically thin emission models with variable element abundances and a photoelectric absorption component. We interpret line flux ratios in the He-like triplet of O\text{vii} as a function of electron density and the UV radiation field. We use the nearby co-eval classical T Tauri star SU Aur as a comparison. AB Aurigae reveals a soft X-ray spectrum, most plasma being concentrated at 1–6 MK. The He-like triplet reveals no signs of increased densities as reported for some accreting T Tau stars in the literature. There are also no clear indications of strong abundance anomalies in the emitting plasma. The light curve displays modulated variability, with a period of \( \approx 42 \) hr. It is unlikely that a nearby, undetected lower-mass companion is the source of the X-rays. Accretion shocks close to the star should be irradiated by the photosphere, leading to alteration in the He-like triplet fluxes of O\text{vii}, which we do not measure. Also, no indications for high densities are found, although the mass accretion rate is presently unknown. Emission from wind shocks is unlikely, given the weak radiation pressure. A possible explanation would be a solar-like magnetic corona. Magnetically confined winds are a very promising alternative. The
X-ray period is indeed close to periods previously measured in optical lines from the wind.
5.1 Introduction

Herbig Ae/Be stars, first defined by Herbig (1960), are young intermediate-mass ($\approx 2-10 \, M_\odot$) stars predominantly located near star-forming dark clouds. They show emission lines in their optical spectra, and their placement in the HR diagram proves that they are pre-main sequence stars (Strom et al. 1972). Herbig stars may therefore be considered to be intermediate-mass analogs of low-mass T Tauri stars (TTS), and in particular of their accreting variant, the classical T Tau stars (CTTS). The analogy between Herbig stars and CTTS extends to infrared excess emission indicative of heated circumstellar material (e.g., disks) and photometric variability. Herbig stars are important in the study of star formation because their evolutionary scenario is intermediate between that of low-mass and high-mass stars. In the former class, the accretion phase, dispersal of the circumstellar disk, and onset of hydrogen burning occur sequentially in time whereas in the latter, the stars enter their main-sequence phase while still being embedded and accreting, making their study much more difficult.

Like their main-sequence descendants, B and A-type stars, Herbig stars are generally supposed to be radiative in their interiors. The lack of a thick convection zone would make the operation of a solar-like $\alpha \Omega$-dynamo impossible, and hence no magnetic fields are expected at the surface of these stars except possibly fossil magnetic fields trapped in the star since the initial cloud contraction phase. However, transient convection may be present in some of these stars during a short phase of deuterium burning in a shell, opening the possibility that some dynamo-generated, non-potential fields develop (Palla & Stahler 1993); alternatively, a dynamo powered by rotational shear energy may generate some surface magnetic fields in rapidly rotating, accreting Herbig stars as well (Tout & Pringle 1995). Although difficult to detect, magnetic fields have recently been measured on several Herbig stars, with (longitudinal) field strengths up to a few 100 G (Donati et al. 1997; Hubrig et al. 2004; Wade et al. 2005). In spectropolarimetry studies of 50 Herbig stars, Wade et al. (2005) detected magnetic fields in 5 of them and discussed these stars to be progenitors of the magnetic Ap/Bp stars. Praderie et al. (1986) have suggested magnetic activity in AB Aur based on their observations of periodically variable blue wings in the Mg II line of this star. Whatever the generation mechanism of magnetic fields, it may be important to include circumstellar disks in the model as the magnetic fields may attach to the inner border of the disk so that they directly interact with accreting material. Moreover, the presence of magnetic fields may also be important in the presence of stellar winds, because the wind could then be magnetically confined and the plasma could be shock-heated to X-ray temperatures (e.g., Babel & Montmerle 1997). We will address this point in Sect. 5.6.3.
X-ray emission is among the best tracers of magnetic fields in stars, although it equally well diagnoses wind or accretion shocks. Damiani et al. (1994) and Zinnecker & Preibisch (1994) were the first to systematically study X-ray emission from Herbig stars with the *Einstein* X-ray observatory and ROSAT, respectively. They reported surprisingly high detection rates of 11/31 and 11/21, respectively. These (and subsequent) studies have investigated X-ray emission in the context of other stellar parameters, with the following principal results: i) The X-ray luminosity, $L_X$, increases with the effective temperature ($T_{\text{eff}}$) and the stellar bolometric luminosity ($L_*$), although the ratio $L_X/L_* \approx 10^{-6} - 10^{-5}$ is higher than in O stars ($L_X/L_* \approx 10^{-7}$) but lower than in T Tau stars ($L_X/L_* \approx 10^{-4} - 10^{-3}$). ii) $L_X$ is not correlated with the projected equatorial velocity $v \sin i$ (Damiani et al. 1994; Zinnecker & Preibisch 1994), thus pointing either to a magnetic dynamo saturation effect or to X-rays not related to dynamo-driven magnetic fields. iii) $L_X$ does correlate with indicators of disks, accretion, and outflows such as infrared excess, mass accretion rate $\dot{M}$, and wind velocity or wind momentum flux. The latter finding may support a non-magnetic origin of the X-rays, for example shocks in unstable winds analogous to more massive O stars (Damiani et al. 1994; Zinnecker & Preibisch 1994).

On the other hand, variability, flares, and extremely high electron temperatures $> 10$ MK clearly favor magnetic processes (Hamaguchi et al. 2000, 2005; Skinner et al. 2004; Giardino et al. 2004; Stelzer et al. 2006) whether based on dynamo-generated or fossil magnetic fields near the star, or star-disk magnetic fields. Surface convection plays an important role in transferring energy into magnetic fields by stirring the magnetic footpoints, a process that leads to coronal heating and mass ejections on the Sun and in magnetically active stars. If the convection zones are shallow as in late-A or early F-type main-sequence stars, then the magnetic dynamos appear to operate rather inefficiently, leading to modest X-ray luminosities of coronal sources that reveal very soft spectra (Panzera et al. 1999). If convection is absent - as in main-sequence A-type stars - then fossil magnetic fields are unlikely to build up non-potential configurations although magnetic activity is widespread among chemically peculiar Bp/Ap stars (Drake et al. 1987, 1994). Alternatively, however, winding-up magnetic fields connecting the star with the inner circumstellar disk may episodically release energy through reconnection, thus heating plasma and possibly ejecting plasmoids that contribute to jets often seen in young stars (Hayashi et al. 1996; Montmerle et al. 2000; Hamaguchi et al. 2005).

Alternatively, X-rays have been suggested to be formed in accretion shocks in the CTTS TW Hya based on its exceptionally soft X-ray spectrum, indications for high electron densities ($\approx 10^{15}$ cm$^{-3}$), and anomalous abundances of N and Ne. Swartz et al. (2005) proposed a similar scenario for the Herbig star HD 163296 based on its unusually soft spectrum ($kT \approx 0.5$ keV).
5.1. Introduction

A caveat is that the majority of Herbig stars are binaries or multiples (Feigelson et al. 2003). A close, unidentified T Tau companion that hides in the strong optical light could easily produce the observed X-rays because X-ray luminosities and electron temperatures of Herbig stars are often quite similar to those of T Tau stars (see for example Stelzer et al. 2006). This emission is commonly interpreted as solar-type coronal magnetic activity. The binary hypothesis has become a relevant model for flaring X-rays in at least two Herbig stars (MWC 297, Hamaguchi et al. 2000; field stars discussed as alternative X-ray sources and companion discovered by Vink et al. 2005; and V892 Tau, Giardino et al. 2004; an 1.5-2\(M_\odot\) companion was discovered by Smith et al. 2005). Stelzer et al. (2006) studied a sample of 17 Herbig stars observed with Chandra and detected 13 of them in X-rays. Of these 13 stars, 7 have a known visual or spectroscopic companion that could be the source of the observed X-rays. Only 35% of the detected sources cannot be explained by known companions. In the latter work, X-ray properties of Herbig stars are found to be very similar to X-rays properties of CTTS, leading to two possibilities: either the mechanism of the X-ray generation is similar for the two types of stars, or the X-rays are generated by (partly unknown) low-mass companions.

We report here the first high-resolution X-ray spectrum of a Herbig star, AB Aurigae, obtained with the Reflection Grating Spectrometer (RGS) on board XMM-Newton. Only the high-resolution spectrum gives access to He-like line triplets, which yield information on densities or UV radiation fields in X-ray emitting regions. This, in turn, constrains the source location and extent (Behar et al. 2004). Further high-resolution spectroscopy enables accurate abundance determination using individual spectral lines (including lines of N and O forming at low temperatures and located at long wavelengths), and, together with CCD spectra obtained with the European Photon Imaging Cameras (EPIC), accurate thermal modeling. Our analysis shows that AB Aur is another very soft source with a moderate X-ray luminosity, but there are no signs of increased electron densities or elevated abundances. We will use the nearby, co-eval, classical T Tau star SU Aur as an ideal comparison star to identify fundamental differences in these X-ray sources with largely differing interiors.

The structure of this chapter is as follows. We describe our target in Sect. 5.2. We then present our data reduction procedures in Sect. 5.3. Sect. 5.4 presents our results from the X-ray spectroscopy, and Sect. 5.5 analyzes information from the He-like triplet of O\(\text{vii}\). We discuss possible models in Sect. 5.6, and conclude in Sect. 5.7.
5.2 AB Aur

Table 5.1 summarizes the basic properties and the principal X-ray parameters of AB Aur. For comparison, the properties of the Herbig star HD 163296, that has been reported to reveal a soft X-ray spectrum (Swartz et al. 2005), and parameters of the CTTS SU Aur are also reported (see also Sect. 5.4 & Sect. 5.6).

New, preliminary estimates of the surface temperature \( T_{\text{eff}} \) and stellar luminosity \( L \) of AB Aur have been derived using the method developed by Fitzpatrick & Massa (1999). By utilizing the available UV (IUE: SWP + LWP) through optical (photometric: UBV) data, the energy distribution can be modeled with surface fluxes represented by Kurucz’s ATLAS13 atmospheric models (Kurucz 1993). See DeWarf et al. (2003) for a more detailed description of this procedure and how it was implemented for SU Aur. A complete description of this analysis, as it pertains to AB Aur, is in preparation (DeWarf et al., in preparation).

The rotation period of AB Aur is controversial. Recent observations suggest an inclination angle \( i \sim 21.5^\circ \) (Corder et al. 2005 and references therein), that would suggest, using \( v \sin i \) and the radius from Table 5.1, a rotation period \( P = 12.9 \) hr. For an extreme value of \( i = 70^\circ \) previously reported for the disk inclination (see references in Corder et al. 2005), we obtain \( P = 33 \) hr. A period of this order is supported by modulations in Mg\( \text{II} \) and Ca\( \text{II} \) lines (Praderie et al. 1986; Catala et al. 1986) and in photospheric lines (Catala et al. 1999) that reveal periods \( P = 32-34 \) hr and \( P = 43-45 \) hr (see Sect. 5.6.1 for a detailed discussion).

To our knowledge, no direct detection of surface magnetic fields of AB Aur has so far been obtained. Catala et al. (1999) estimated an upper limit to the strength of the magnetic field of 300 G.

5.3 Observations and Data Analysis

AB Aur was detected in an XMM-Newton (Jansen et al. 2001) observation pointing at the nearby CTTS SU Aur (separation between SU Aur and AB Aur: \( \approx 2.5' \)). The observation was retrieved from the archive as part of the XMM-Newton Extended Survey of Taurus Molecular Cloud (XEST) described in Güdel et al. (2007a) (XEST observation number of AB Aur: XEST-26-043, and of SU Aur: XEST-26-067). Table 5.2 summarizes the observing parameters. The two European Photon Imaging Cameras (EPICs) of the MOS type (Turner et al. 2001) and the Reflection Grating Spectrometers (RGSs; den Herder et al. 2001) were active during the observation, while the EPIC PN camera was out of operation. Both MOS instruments observed in full frame mode and used the thick filter to suppress excessive optical load from
Table 5.1: Parameters for AB Aur, SU Aur, and HD 163296

<table>
<thead>
<tr>
<th>Parameter</th>
<th>AB Aur</th>
<th>HD 163296</th>
<th>SU Aur</th>
</tr>
</thead>
<tbody>
<tr>
<td>Magnetic field</td>
<td>Y? (1)</td>
<td>Y? (15)</td>
<td>Y</td>
</tr>
<tr>
<td>Spectrum</td>
<td>B9.5e-A0 (2,3)</td>
<td>A1Ve (4)</td>
<td>G2 (5)</td>
</tr>
<tr>
<td>$L_*$ [$L_\odot$]</td>
<td>49 (14)</td>
<td>30 (4)</td>
<td>9.9 (6)</td>
</tr>
<tr>
<td>Mass [$M_\odot$]</td>
<td>2.7 (14)</td>
<td>2.3 (4)</td>
<td>1.9 (X)</td>
</tr>
<tr>
<td>Radius [$R_\odot$]</td>
<td>2.3 (14)</td>
<td>2.1 (4)</td>
<td>3.1 (X)</td>
</tr>
<tr>
<td>$A_V$ [mag]</td>
<td>0.25 (21)</td>
<td>0.25 (4)</td>
<td>0.5 (7)</td>
</tr>
<tr>
<td>$T_{\text{eff}}$ [K]</td>
<td>9750 (8)</td>
<td>9300 (4)</td>
<td>5860 (6)</td>
</tr>
<tr>
<td>Age [Myr]</td>
<td>4 (14)</td>
<td>4 (9)</td>
<td>4 (7)</td>
</tr>
<tr>
<td>Disk</td>
<td>Y (19)</td>
<td>450 AU (9)</td>
<td>Y</td>
</tr>
<tr>
<td>Companions?</td>
<td>evidence (18)</td>
<td>not det. (9)</td>
<td>...</td>
</tr>
<tr>
<td>Radio spectrum</td>
<td>wind (11)</td>
<td>wind (12)</td>
<td>...</td>
</tr>
<tr>
<td>EW(Hα) [Å]</td>
<td>27 (13)</td>
<td>2-6 (X)</td>
<td></td>
</tr>
<tr>
<td>vsini [km s$^{-1}$]</td>
<td>80 (16)</td>
<td>120 (17)</td>
<td>65 (20)</td>
</tr>
<tr>
<td>$kT$ [keV]</td>
<td>0.46 (23)</td>
<td>0.49 (9)</td>
<td>1.9 (23)</td>
</tr>
<tr>
<td>log $L_X$ [erg s$^{-1}$]</td>
<td>&lt;30.3 (3)</td>
<td>29.6 (9)</td>
<td>30.9 (23)</td>
</tr>
<tr>
<td></td>
<td>29.5 (10)</td>
<td></td>
<td></td>
</tr>
<tr>
<td></td>
<td>29.6 (23)</td>
<td></td>
<td></td>
</tr>
<tr>
<td>log $L_X/L_*$</td>
<td>-5.8 (10)</td>
<td>-5.48 (9)</td>
<td>-3.7 (23)</td>
</tr>
<tr>
<td></td>
<td>-5.6 (23)</td>
<td></td>
<td></td>
</tr>
<tr>
<td>log $N_H$ [cm$^{-2}$]</td>
<td>20.7 (23, X)</td>
<td>20.88 (9)</td>
<td>21.4 (23)</td>
</tr>
<tr>
<td>X-ray variable?</td>
<td>slow (23)</td>
<td>marginal? (9)</td>
<td>flares (22)</td>
</tr>
</tbody>
</table>

REFERENCES given in parentheses: 1 Praderie et al. (1986) - see text for further discussion; magnetic field assumed for CTTS. 2 Hamann & Persson (1992); 3 Hamaguchi et al. (2005) (and references therein); 4 van den Ancker et al. (1998); 5 Kenyon & Hartmann (1995); 6 Luhman (2004); 7 DeWarf et al. (2003); 8 Acke & Waelkens (2004); 9 Swartz et al. (2005) (and references therein); 10 Zinnecker & Preibisch (1994); 11 Skinner et al. (1993); 12 Blondel et al. (1993) 13 Herbig & Bell (1988); 14 deWarf & Fitzpatrick, private communication; 15 Deleuil et al. (2005); 16 Böhm & Catala (1993); 17 Hillenbrand et al. (1992); 18 Baines et al. (2006); 19 Corder et al. (2005); 20 Hartmann & Stauffer (1998); 21 Roberge et al. (2001); 22 Robrade & Schmitt (2006); 23 this work. X see XEST survey results and tables in Güdel et al. (2007a) for derivations and references.
AB Aur. The EPIC detectors operate in the energy range of 0.15–15.0 keV with a medium spectral resolution of approximately $E/\Delta E = 20 - 50$. The RGSs are suited for high-resolution spectroscopy in the wavelength range of 6–35 Å and have a resolution of $\Delta \lambda \approx 60 - 76$ mÅ. The RGS detectors contained the dispersed spectra of both AB Aur and SU Aur with sufficient separation to make mutual spectral contamination negligible (see below). The spectrum of SU Aur has recently been discussed by Robrade & Schmitt (2006).

The data were reduced using the Science Analysis System (SAS) version 6.1. The EPIC MOS data were reduced using the task `emchain` and the sources were detected using the maximum likelihood detection algorithm `emldetect` (see Güdel et al. 2007a for further details). To extract the two spectra from each RGS detector, we proceeded as follows. We first applied the standard processing performed by the RGS metatask `rgsproc` to each source position. We extracted the total (source+background) spectra and the background spectra separately. For both SU Aur and AB Aur, we included 85% of the cross-dispersion Point Spread Function (PSF) as also done in Chapter 3 ($xpsfinc1 = 85$). For the background, the exclusion region with regard to the cross-dispersion PSF and the inclusion region of the pulse-height distribution were kept at default values (95% of the PSF and 90% of the pulse-height distribution, respectively, i.e., $xpsfexcl = 95$ and $pdistincl = 90$). We verified that this way the spatial extraction regions on the detector were adjacent to each other but not overlapping. This means that each spectrum collects approximately $(100-85)/2 = 7.5\%$ of the counts of the other spectrum. The contamination is such that counts from SU Aur are shifted in wavelength by approximately -0.4 Å in the AB Aur spectrum within the wavelength range of interest here (12–22 Å), and contaminating counts from AB Aur are shifted by +0.4 Å in the SU Aur spectrum.

The background defined for each source outside its source extraction region is now still contaminated by the other source. Therefore, we defined the coordinate of the secondary source (AB Aur for the SU Aur spectrum and vice versa), added them to the source list using the `rgssources` task, and computed

### Table 5.2: Observing log

<table>
<thead>
<tr>
<th>XMM-Newton ObsID</th>
<th>0101440801</th>
</tr>
</thead>
<tbody>
<tr>
<td>Boresight RA (J2000.0)</td>
<td>4h 55m 59.0s</td>
</tr>
<tr>
<td>Boresight δ (J2000.0)</td>
<td>30° 34′ 02″</td>
</tr>
<tr>
<td>Start time (UT)</td>
<td>2001-09-21 01:34:17</td>
</tr>
<tr>
<td>Stop time (UT)</td>
<td>2001-09-22 13:34:31</td>
</tr>
<tr>
<td>Exposure time (s)</td>
<td>129614</td>
</tr>
<tr>
<td>MOS Mode and Filter</td>
<td>Full Frame, thick filter</td>
</tr>
</tbody>
</table>
5.3. Observations and Data Analysis

Table 5.3: Wavelength ranges used for the spectral fitting

<table>
<thead>
<tr>
<th>Instrument</th>
<th>AB Aur</th>
<th>SU Aur</th>
</tr>
</thead>
<tbody>
<tr>
<td>MOS1</td>
<td>–</td>
<td>1.5–9.35 Å</td>
</tr>
<tr>
<td>MOS2</td>
<td>1.5–9.35 Å</td>
<td>–</td>
</tr>
<tr>
<td>RGS1</td>
<td>10.0–28.0 Å</td>
<td>8.3–25.0 Å</td>
</tr>
<tr>
<td>RGS2</td>
<td>8.3–26.5 Å</td>
<td>8.3–25.0 Å</td>
</tr>
</tbody>
</table>

A new extraction map, with the secondary source excluded from the background, using the *rgsregion* task. Finally, we extracted each spectrum again, using only the regions outside the SU Aur and AB Aur source regions for the background spectra. The background for the AB Aur spectrum is extracted outside the two adjacent source regions. The background portion on the far side of the SU Aur spectrum again comprises approximately 7.5% of the SU Aur source counts. Because this background spectrum is subtracted from the total spectrum extracted at the AB Aur position, the contamination is approximately corrected for. Further considering the low S/N, the similar line fluxes in both spectra, and the line-dominated spectrum of AB Aur (see below, while most of the counts in SU Aur are in continuum and therefore distributed in wavelength), the mutual contamination is negligible and much below the noise level. For example, no significant mutual contamination is seen by the strong O\textsuperscript{viii} lines above the noise level in either of the spectra (see presentation of the RGS spectra in Fig. 5.3 below). The periods affected by high background flaring were excluded from the spectral analysis.

Because we wanted to put weight on the high-resolution spectra obtained with RGS, we performed the data analysis on both RGS spectra, adding only the short-wavelength portion of one of the two MOS detectors in order to access line features of Mg, Si and Fe that are important for our abundance analysis (see Chapter 2). As in Chapter 3, we used MOS1 for SU Aur for this purpose, while for AB Aur, we preferred MOS2 because an unidentified background feature distorted the MOS1 spectrum around a wavelength of 9.3 Å (not present in MOS2). The wavelength intervals of each instrument used for the data analysis are summarized in Table 5.3. As a check, we confirmed our results by using the entire useful energy range of MOS (0.2–10 keV), and the results are consistent with those reported here.

We fitted the spectra in XSPEC (Arnaud 1996) using optically-thin collisionally-ionized plasma models calculated with the Astrophysical Plasma Emission Code (APEC; Smith et al. 2001). In order to account for calibration discrepancies between the RGS and the MOS instruments, we added constants as effective-area renormalization factors. These factors were fixed at 1.0 for MOS and at 1.05 for both RGS (see Kirsch et al. 2004).
Since we used the $\chi^2$-statistic for our spectral fitting, we binned the spectra to a minimum of 20 counts per bin for the RGS and a minimum of 15 counts per bin for the MOS. The resulting bin width varies between 0.04 Å (in the O vii line) and 0.73 Å (just longward of the O viii Lyα line) in the RGS, and between 0.15 Å and 1.4 Å in the MOS spectrum.

For AB Aur the hydrogen column density $N_H$ was fixed at the value found in the XEST survey analysis, namely $5 \times 10^{20}$ cm$^{-2}$, which is consistent with the recent measurements of $A_V = 0.25$ mag reported by Roberge et al. (2001), assuming a standard conversion $N_H \approx 2.0 \times 10^{21}$ $A_V$ applicable to the interstellar medium (Vuong et al. 2003, and references therein). This value agrees quite well with the $N_H$ found explicitly by Roberge et al. (2001), $N_H = 4.4 \times 10^{20}$ cm$^{-2}$. For SU Aur, we needed to fit $N_H$ in order to get a good fit to the RGS spectra.

The spectra were fitted with two different models: a model describing a differential emission measure distribution (EMD) approximated by two power laws as used in the XEST survey analysis (Güdel et al. 2007a; Chapter 1, Sect. 1.6), and a model with two or three isothermal plasma components. For both models we applied a procedure in which we simultaneously fitted large wavelength intervals of the three spectra with template spectra computed in XSPEC. Alternative, iterative methods based on extracted line fluxes are not feasible given the low S/N ratio of our spectra. There are only a few explicitly measurable lines, and each set of lines from a given element is confined to a narrow formation temperature interval. Also, such methods cannot be applied to MOS CCD data. Previous studies have shown excellent agreement between iterative methods and the method applied here (see Chapter 2). The EMD model is approximated by a grid of isothermal components spaced regularly by 0.1 dex in $\log T$ in such a way that the lower-$T$ and the higher-$T$ portions are each described by a power law. This model has been suggested from our previous work on high-resolution X-ray spectroscopy of young solar analogs (Chapter 2), and is described by

$$Q(T) = \begin{cases} 
EM_0 \cdot \left(\frac{T}{T_0}\right)^\alpha, & \text{for } T \leq T_0 \\
EM_0 \cdot \left(\frac{T}{T_0}\right)^\beta, & \text{for } T > T_0
\end{cases} \quad (5.1)$$

where $T_0$ is the temperature where the two power-laws cross, and $EM_0$ is the EM per $\log T$ at this crossing point. The slopes of the power laws below and above $T_0$ are $\alpha$ and $\beta$, respectively. The power laws have high-energy and low-energy cut-offs at $\log T = 8$ and $\log T = 6$, respectively. Given the poor energy resolution of CCDs below 1 keV, we fixed $\alpha$ at a value of +2 throughout the XEST survey, and we will do so here for comparison. A slope of 2 is compatible with findings from DEM analyses in other pre-main sequence and main-sequence active stars (Chapter 2; Argiroffi et al. 2004).

However, because the X-ray emission of a Herbig Ae/Be star could be non-
coronal, or the coronal thermal structure could be unlike that in late-type T Tau stars, it is possible that the simplified EMD model proposed here is inappropriate. In a separate approach, we will therefore also fit $\alpha$. Further, because the power-law EMD structure might be inappropriate, we will test our results using a multi-temperature model that can reveal the temperatures where most EM resides. We therefore analyzed the AB Aur spectra with a 2-component model. A third component was not needed, given the low signal-to-noise ratio of the spectrum and, as we will describe below, its rather narrow range of temperatures in which emitting plasma is found. Because spectral lines are formed over an interval of typically 0.3 dex in temperature, a temperature range of 2-7 MK (the range of formation temperatures present in the spectrum) will require no more than two thermal components. We also analyzed the SU Aur spectra with a 3-component model.

For both approaches, we fitted the abundances of the lines observed in the spectra simultaneously with the thermal models. Abundances that do not show significant features in the spectra, such as those of C, S, Ar, Ca and Al (and N for SU Aur), were fixed at the values used in the general XEST survey data analysis (C=0.45, N=0.788, S= 0.417, Ar=0.55, Ca=0.195, and Al=0.5, Güdel et al. 2007a). These abundances were arranged in such a way that they describe a weak “inverse First Ionization Potential Effect” (higher FIP implies higher abundances relative to the photospheric values, the latter assumed to be solar), as often observed in young stars.

Finally, we computed the X-ray luminosities, $L_X$, from the spectral model in the energy range of 0.3–10.0 keV, assuming a distance of 140 pc (DeWarf et al. 2003 and references therein).

### 5.4 Results

#### 5.4.1 Light Curves

We first present the X-ray light curve of AB Aur in Fig. 5.1. We have co-added counts from both EPIC MOS detectors after background subtraction. The background contribution to the total light curve was in fact negligible except during the final 5 ks when a background flare occurred that reached $\approx 50 \%$ of the total count rate. The observed variability can therefore not be attributed to imperfect background treatment but is intrinsic to the stellar source. We see slow variability on time scales of about one day, by somewhat less than a factor of two. In the variability analysis of the XEST sources presented by Stelzer et al. (2007), the variability of AB Aur was confirmed by the Kolmogorov-Smirnov test. No significant trend is seen in the hardness, defined as the ratio between the count rates in the hard (1-4 keV) and the soft (0.3-1 keV) band. We have fitted the light curve with a sine function which is
also plotted in Fig. 5.1. The fit is excellent ($\chi^2 = 12.9$ for 22 dof) with a period of 42.2$^{+4.4}_{-3.7}$ hr in the total band (in the soft and hard band we find 48.5$^{+11.6}_{-7.8}$ hr and 40.4$^{+5.2}_{-3.2}$ hr, respectively). A similar modulation of $\approx 42 - 45$ hr was found previously for the HeI and MgII lines (see discussion below).

We have studied the light curve variability with a statistical test to assess the presence of a flare in the last 30 ks, where an increase in the count rates can be seen, particularly in the hard and total light curves (second and third panel in Fig. 5.1, respectively). We first tested our data against a constant count-rate model using the $\chi^2$-statistic. We find the probability for this part of the light curve to be constant, $P$(const)=0.73 for the total band, and $P$(const)=0.30 for the hard band. As a second test, we used the Kolmogorov-Smirnov statistic, to obtain $P$(const)=0.09 and $P$(const)=0.07 for the total and hard light curves, respectively. The variability of the last 30 ks in the light curves is thus at best marginal.
5.4. Results

5.4.2 Spectra

Fig. 5.2 compares the background-subtracted EPIC MOS spectra of AB Aur and SU Aur. A number of differences are obvious. The spectrum of SU Aur reveals signatures of a very hot coronal plasma, with outstanding lines of Mg xi, Mg xii, Si xiii and, most notably, the Fe K complex mostly due to Fe xxv at 6.7 keV. A steep drop toward the lowest energies indicates considerable photoelectric absorption. In contrast, the AB Aur spectrum falls off rapidly above 1 keV, showing its peak flux around 0.7-0.9 keV. This flux peak is mainly due to lines of Fe xvii. These spectral properties let us anticipate a soft source for AB Aur, and the shallow fall-off toward low energies indicates rather low photoelectric absorption.

Fig. 5.3 shows fluxed, co-added RGS1+2 spectra of SU Aur (top) and AB Aur (bottom) in the line dominated region. The spectra have been rebinned to a bin width of 0.042 Å for AB Aur and 0.035 Å for SU Aur. These spectra further corroborate the differences between the two X-ray sources. While SU Aur reveals a strong continuum, indicating a hot source, there is little evidence for continuum emission in AB Aur. Further, the flux ratio between the Ne x Lyα line at 12.1 Å (formation temperature 6.3 MK) and the Ne ix resonance line at 13.44 Å (formation temperature 4 MK) is considerably higher in SU Aur than in AB Aur, again emphasizing the dominance of hot plasma in the former. The Ne ix line feature at 13.5 Å in SU Aur is dominated by Fe xix, which is formed at higher temperatures (formation temperature 7.9 MK). Two further features are striking: First, the Fe lines of SU Aur are very strong,
dominating the spectrum and comparable in flux with the O\text{VIII} Ly\alpha and the Ne\text{X} Ly\alpha lines (although the O\text{VIII} line is partly suppressed by photoelectric absorption). Such line ratios are unusual among very active, main-sequence solar analogs where Fe line fluxes are modest due to low Fe abundance (see Chapter 2). We thus anticipate an unusually high abundance of Fe in this spectrum.

The second feature of interest here is the unusually high flux of the O\text{VII} lines of AB Aur, despite some photoelectric absorption that suppresses the flux at these wavelengths. The total O\text{VII} flux appears to be similar to the flux in the O\text{VIII} Ly\alpha line, which is a property of very inactive stellar coronae with temperatures of 3–5 MK (Chapter 2). The feature at 18.5 Å is neither coincident with the O\text{VII} He\beta line nor due to contamination by the O\text{VIII} Ly\alpha line of SU Aur. Those two line features would both be slightly but significantly longward of this wavelength, namely at 18.6 Å. The 18.5 Å feature is due to a 3σ spike exclusively in the RGS2 detector and is therefore spurious. The O\text{VII} lines are not present in the spectrum of SU Aur due to considerable photoelectric absorption.

5.4.3 Thermal Structure

We now present the numerical results of our spectral fits to these data. Table 5.4 lists the numerical results for both models and both stars. In Figure 5.4 we plot the data together with the best fit of the EMD model for AB Aur.\footnote{The low bin at 16 Å in RGS2, coincident with the O\text{VIII}+Fe\text{XVIII} lines, corresponds to a CCD gap. At the same time, somewhat increased flux in the RGS1 spectrum slightly longward of the two lines, probably due to imperfect background subtraction, increases the RGS1 flux in the respective bin in Fig. 5.4 above our fit. We checked at higher-resolution...}
5.4. Results

Figure 5.4: Data and fitted spectrum of AB Aur (EMD model). The best-fit model is shown by the histograms in the wavelength region used for the fit. For plotting purposes, the MOS2 spectrum has been shifted along the y-axis by 0.001 cts s$^{-1}$ Å$^{-1}$.

We define the average coronal temperature, $T_{av}$, as the logarithmic average of all temperatures used in the model, applying the corresponding EM as weights. This measure corresponds to the electron temperature itself for an isothermal plasma. For a continuous EMD as discussed here, $T_{av}$ represents a temperature grossly characteristic of the spectral shape. This is also true for multi-component plasma (e.g., 2-$T$ or 3-$T$ plasmas) although there might be no plasma present at $T_{av}$ itself. This also occurs when 1-$T$ fits are made to low-quality coronal spectra.

Both the multi-temperature and the EMD models provide similar results: $T_{av}$ and the X-ray luminosities are similar. The 2$T$ or 3$T$-model fits and the EMD model fits show similar $\chi^2$ values, and are therefore of statistically equal quality. Most of the abundance values agree between the two methods, confirming the robustness of our results.

that the O VIII+Fe XVIII line is indeed correctly fitted.
Table 5.4: Results of spectral model interpretation

<table>
<thead>
<tr>
<th></th>
<th>AB Aur</th>
<th>SU Aur</th>
</tr>
</thead>
<tbody>
<tr>
<td></td>
<td>2T model</td>
<td>DEM model</td>
</tr>
<tr>
<td></td>
<td>= 0.5</td>
<td>= 0.5</td>
</tr>
<tr>
<td>$N_H$ [10^{21} cm^{-2}]</td>
<td>0.45 (0.22, 0.78)</td>
<td>0.57 (0.30, 1.13)</td>
</tr>
<tr>
<td></td>
<td>0.20 (0.16, 0.42)</td>
<td>0.22 (0.13, 0.32)</td>
</tr>
<tr>
<td></td>
<td>0.60 (0.41, 0.88)</td>
<td>0.62 (0.48, 1.04)</td>
</tr>
<tr>
<td></td>
<td>0.21 (0.09, 0.38)</td>
<td>0.28 (0.13, 0.74)</td>
</tr>
<tr>
<td></td>
<td>0.70 (0.48, 1.01)</td>
<td>0.90 (0.60, 1.32)</td>
</tr>
<tr>
<td>$T_1$ [MK]</td>
<td>2.45 (2.10, 2.81)</td>
<td>–</td>
</tr>
<tr>
<td>$T_2$ [MK]</td>
<td>6.99 (6.62, 7.41)</td>
<td>–</td>
</tr>
<tr>
<td>$T_3$ [MK]</td>
<td>–</td>
<td>–</td>
</tr>
<tr>
<td>$E_{M1}$ [10^{52} cm^{-3}]</td>
<td>2.11 (0.87, 3.67)</td>
<td>–</td>
</tr>
<tr>
<td>$E_{M2}$ [10^{52} cm^{-3}]</td>
<td>3.44 (3.06, 4.31)</td>
<td>–</td>
</tr>
<tr>
<td>$E_{M3}$ [10^{52} cm^{-3}]</td>
<td>–</td>
<td>–</td>
</tr>
<tr>
<td>$T_{0}$ [MK]</td>
<td>–</td>
<td>4.38 (2.69, 5.68)</td>
</tr>
<tr>
<td>$T_{av}$ [MK]</td>
<td>–</td>
<td>5.47 (0.00, 1.20)</td>
</tr>
<tr>
<td>$\alpha$</td>
<td>–</td>
<td>2.00</td>
</tr>
<tr>
<td>$\beta$</td>
<td>–</td>
<td>-1.9 (-2.57, -1.52)</td>
</tr>
<tr>
<td>$EM$ [10^{52} cm^{-3}]</td>
<td>4.69</td>
<td>5.21</td>
</tr>
<tr>
<td>$L_X$ [10^{30} erg/s](^{\dagger})</td>
<td>0.40</td>
<td>0.39</td>
</tr>
<tr>
<td>$L_X$ [10^{30} erg/s](^{\ddagger})</td>
<td>0.55</td>
<td>0.53</td>
</tr>
<tr>
<td>$dof$</td>
<td>78</td>
<td>79</td>
</tr>
</tbody>
</table>

Notes:

1. 68% error ranges are given in parentheses.
2. Fitted with $\alpha$ held fixed at 2 like in the XEST survey (Güdel et al. 2007a).
3. Element abundances are with respect to solar values given by (Anders & Grevesse 1989).
4. Total EM integrated over temperature bins between log $T[K]=6$ – 7.9 (see Güdel et al. 2007a for more details).
5. Determined in the 0.3-10.0 keV band.
6. Determined in the 0.1-10.0 keV band.
For AB Aur we report the results for EMD models with $\alpha$ fixed at 2 and with $\alpha$ free (second and third columns of Table 5.4, respectively). The peak temperature $T_0$ increases from 4.4 MK for $\alpha$ fixed at 2 to 6.3 MK for free $\alpha$. However, the main results are in very close agreement for the two approaches: the abundances are consistent within the error bars, the X-ray luminosities are essentially unchanged, and $T_{\text{av}}$ is only slightly smaller when we fit $\alpha$. This suggests that the two solutions are equivalent. On the other hand, the errors are larger if $\alpha$ is a free fit parameter (except for the Mg abundance). Therefore, and for consistency with the fits of the EPIC spectra in the XEST survey, we use the results from the EMD fit with $\alpha$ fixed at a value of 2.

The $N_H$ value for SU Aur was found to be $N_H = (3.1 - 3.2) \times 10^{21}$ cm$^{-2}$, which agrees well with the result reported by Skinner & Walter (1998) from an ASCA observation, $N_H = 2.8 \times 10^{21}$ cm$^{-2}$ and is only somewhat higher than expected from $A_V = 0.9$ mag (Kenyon & Hartmann 1995).

The results confirm the peculiar thermal structure of the AB Aur source. The average temperature from the 2-$T$ fit is 4.7 MK, while the EMD model peaks at 4.4 MK and rapidly falls off toward higher temperatures, with a power-law slope of -1.9. Such low temperatures are unusual for coronae of young stars where usually temperatures in excess of 10 MK are found. An example is SU Aur: The 3-$T$ model shows the largest amount of EM at the highest temperature at $\approx 40$ MK, with $T_{\text{av}} \approx 23$ MK. Again, the EMD model supports this finding, where we find $T_0$ at 7.7 MK beyond which the EMD is nearly flat, resulting in an average temperature of 20 MK. Franciosini et al. (2006, in preparation) have analyzed the same RGS spectrum of SU Aur using a line-based analysis to derive the EMD from the measurement of individual line fluxes. Their results are consistent with ours: the EMD peaks at $T = 10$ MK, with an indication of a significant amount of material above $\sim 20$ MK; however, below the peak they find a steeper slope ($\sim 3$) than adopted here.

5.4.4 Abundances

The element abundances in the X-ray sources are plotted in Fig. 5.5. The filled circles designate abundance ratios relative to the solar photospheric mix (Anders & Grevesse 1989; Grevesse & Sauval 1999 for Fe). The abundances of AB Aur show a nearly flat distribution; neither a strong First Ionization Potential (FIP) effect (i.e., overabundant low-FIP elements) nor a strong inverse FIP effect (i.e., increasing abundances with increasing FIP) is seen, in contrast to the usual findings in young, active stars (e.g., Argiroffi et al. 2004, Chapter 3). Because the emitting material originates, if located in a corona or a stellar wind, in the stellar photosphere, a comparison of the X-ray derived abundances with photospheric values will be important. Fortunately, a few photospheric abundances have been measured for AB Aur (Acke & Waelkens 2004), and the resulting normalized abundance ratios are shown by the open
circles. Acke & Waelkens (2004) report a particularly low photospheric abundance of Fe (≈ 13% of the solar value given by Grevesse & Sauval 1999). The renormalized abundance distribution is still ambiguous: the coronal abundance of Fe is higher than the photospheric value, but a clear FIP dependent abundance distribution is not visible.

The derived abundances in the SU Aur source are atypical for young, active stars and T Tau stars. Their X-ray sources usually show a well-expressed inverse FIP effect (e.g., Chapter 3; Argiroffi et al. 2004). In contrast, SU Aur shows O and Ne abundances clearly lower than Fe, which is a defining signature of a solar-type FIP effect. As suggested earlier, the absolute Fe abundance is quite high (0.67-0.81 times the solar photospheric value). Such high values are usually reported for relatively inactive stars, while magnetically active stars reveal much greater depletion of Fe (Güdel 2004a; see also Chapter 2).

5.4.5 Luminosities

The only previous report on X-ray emission from AB Aur that we are aware of is by Zinnecker & Preibisch (1994). This allows us to study possible long-term changes in the X-ray output. Zinnecker & Preibisch (1994) give an X-ray luminosity of \((0.3 \pm 0.09) \times 10^{30} \text{ erg s}^{-1}\) in the energy range 0.1-2.4 keV. In contrast, we find \(L_X \approx 0.4 \times 10^{30} \text{ erg s}^{-1}\) in the energy range 0.3-10.0 keV, and

---

2The photospheric abundances of AB Aur are, with respect to the solar photosphere (error ranges in parentheses): N: 1.38; O: 1.62 (1.38-1.91); Si: 0.78 (0.72-0.83); Fe: 0.13 (0.10-0.18), referring to the solar abundances of Anders & Grevesse (1989) for N, O, and Si, and those of Grevesse & Sauval (1999) for Fe.
$L_X \approx (0.5 - 0.6) \times 10^{30}$ erg s$^{-1}$ in the 0.1–2.4 keV range, that is, almost twice as much as Zinnecker & Preibisch (1994). We note, however, that Zinnecker & Preibisch (1994) estimated $L_X$ directly from count rates, using a conversion factor applicable for a temperature of 1 keV and an $N_H$ corresponding to $A_V = 0.65$ mag. Modeling the ROSAT count rate for $N_H, kT$ and abundances found from our spectral analysis still results in only $(0.3 - 0.4) \times 10^{30}$ erg s$^{-1}$ in the 0.1–2.4 keV range. AB Aur was in a more active state during our observation, but this source is slowly variable on time scales of hours (Fig. 5.1). Our luminosity of SU Aur is $L_X \approx (7.4 - 7.8) \times 10^{30}$ erg s$^{-1}$ in the 0.3–10 keV range. This compares well with values reported by Skinner & Walter (1998), i.e., $(8.4 \pm 0.09) \times 10^{30}$ erg s$^{-1}$ in the energy range 0.5–10.0 keV.

5.4.6 Variability of Spectral Fit Parameters

In view of the sinusoidal variation of the X-ray count rate in Fig. 5.1, we tested which spectral fit parameters are mainly responsible for the modulation. We restricted this study to the combined EPIC MOS1+2 data because the RGS signal becomes too weak if the data are split. We first fitted the entire MOS data set with two thermal components to obtain a fit very similar to the one reported in Table 5.4, apart from some deviations in the element abundances (most of which are difficult to derive from MOS, in particular O and Ne, given the modest resolution and severe blending). We then split the observation into “high state” (first half) and “low state” (second half). Starting with the above model fit, we tested whether i) an adjustment of $N_H$ (variation due to selective absorption), ii) a renormalization of the emission measures by a common factor (variation of $L_X$), or iii) some changes in all of $N_H, kT$, and EM are required. A variation of $N_H$ could clearly be excluded. This is supported by Fig. 5.1 that shows that the hard photons vary in concert with the soft photons, while they are not significantly affected by the weak photoelectric absorption. To quantify this, we measured the amplitude of the sine function relative to the “zero level” in each light curve and found them to be the same in each energy range, within the errors. No significant changes in $kT$ were required, while a simple renormalization of the EMs by a common factor produced perfect fits. We conclude that the variation of the X-rays are either due to an intrinsic change in the luminosity with other plasma parameters remaining equal, or due to an energy-independent filtering of photons (e.g., due to partial eclipses by the star).
5.5 He-like Triplets, Densities and the Radiation Field

We now discuss the helium-like line triplet of O\textsc{vii} for AB Aur. The flux ratio between the forbidden and the intercombination lines at 22.1 Å and 21.8 Å, respectively, is density-sensitive roughly in the range of electron densities between $10^{10}$ cm$^{-3}$ and $10^{12}$ cm$^{-3}$ (Gabriel & Jordan 1969) for the following reason: if the electron collision rate is sufficiently high, electrons in the upper level of the forbidden transition, $1s2s\ ^3S_1$, do not return to the ground level, $1s^2\ ^1S_0$, instead they are collisionally excited to the upper levels of the intercombination transitions, $1s2p\ ^3P_{1,2}$, from where they decay radiatively to the ground state. They thus enhance the flux in the intercombination line and weaken the flux in the forbidden line. However, photons in a UV radiation field may excite the same transition. The relevant photon energies correspond to the energy difference of the two upper states, and this corresponds to a wavelength of 1630 Å for the O\textsc{vii} triplet. The UV radiation field is thus important for stars with $T_{\text{eff}}$ of about $10^4$ K and more. Because $T_{\text{eff}}$ of AB Aur has been quoted to be around $10^4$ K (Table 5.1), we need to consider the radiation term. We follow Blumenthal et al. (1972) for a rough estimate. The measured ratio $R = f/i$ of the forbidden to the intercombination line flux can be written as

$$R = \frac{R_0}{1 + \phi/\phi_c + n_e/N_c} = \frac{f}{i}$$  \hspace{1cm} (5.2)

where $R_0$ is the limiting flux ratio at low densities ($R_0 \approx 3.85$ for O\textsc{vii} at the maximum formation temperature), $N_c$ is the critical density at which $R$ drops to $R_0/2$ ($N_c \approx 3.4 \times 10^{10}$ cm$^{-3}$ for O\textsc{vii}), and $\phi_c$ is the critical photoexcitation rate. The influence of the radiation field is given by

$$\phi = \frac{3(1 + F)c^3}{8\pi h\nu^3} \frac{A(1s2p\ ^3P \rightarrow 1s2s\ ^3S_1)}{A(1s2s\ ^3S_1 \rightarrow 1s^2\ ^1S_0)} u_\nu$$  \hspace{1cm} (5.3)

where the radiation field energy density is given by Planck’s equation,

$$u_\nu = \frac{W}{c^3} \frac{1}{\exp(h\nu/kT_{\text{eff}}) - 1}$$  \hspace{1cm} (5.4)

in which we have introduced a geometric dilution factor (e.g., Mewe & Schrijver 1978),

$$W = \frac{1}{2} \left(1 - \left[1 - \left(\frac{R}{d}\right)^2\right]^{1/2}\right)$$  \hspace{1cm} (5.5)

where $R$ is the stellar radius and $d$ is the distance of the source from the center of the star (see also Ness et al. 2001 for a more detailed presentation). At the
surface, $W = 0.5$. Further, in Eq. (5.3), the $A$ terms are the spontaneous transition probabilities, $(8.12 \times 10^7 \text{ s}^{-1}$ and $1.04 \times 10^3 \text{ s}^{-1}$, respectively, after Blumenthal et al. 1972) and $F$ is an expression that we approximate by adopting the maximum line formation temperature for O vii ($\approx 2 \text{ MK}$), which yields $F = 0.42$ for this ion (Blumenthal et al. 1972). We have also assumed that the radiation field exactly corresponds to the stellar photospheric $T_{\text{eff}}$.

In a thermal plasma, the flux of the resonance line, $r$, is larger than the flux of the forbidden line, $f$, and under conditions relevant to us, namely $T > 1.5 \text{ MK}$, also the sum of $f+i$ is smaller than $r$, $G = (f+i)/r < 1$ (the “$G$ ratio”, see, e.g., Porquet et al. 2001). The considerable errors in our measurements make the extraction of individual lines and the separate treatment of $R$ and $G$ problematic. To obtain self-consistent $R$ and $G$ while fulfilling other conditions from the spectral fit to the entire spectrum, we proceeded as follows. We adopted the optimum parameters from the 2-$T$ fit and kept all parameters fixed, except for the electron density and (for slight adjustments to the total O vii line flux) the emission measure of the cooler component. The thermal structure thus sets a requirement on $G$ (Porquet et al. 2001) and also fixes the faint continuum required by other lines. We then performed a fit only to the spectral range around the O vii triplet, in the wavelength interval 21.4–22.4 Å.

We first assume a negligible influence of the radiation field, i.e., we assume that $f/i$ is controlled by the electron density. Our fit procedure makes use of the implementation of the He-like triplet calculations in XSPEC’s vmekal code. Varying the electron density changes both the $i$ and $f$ flux until a best fit is obtained. We performed this procedure for various data binning schemes, using bin widths of 45 mÅ or 56 mÅ. Each time, the best-fit parameters converged to values below the low-density limit for O vii ($n_e < 10^{10} \text{ cm}^{-3}$, corresponding to $R = R_0$). We then varied $n_e$ to find the 68% and 90% upper limits. These themselves fluctuated for different binning schemes; for the average 90% upper limit, we find $n_{e,\text{max},90} \approx (1.3 \pm 0.4) \times 10^{11} \text{ cm}^{-3}$, corresponding to the measured $f/i = 0.95$. For the 68% errors, we find $n_{e,\text{max},68} \approx (4.2 \pm 1.2) \times 10^{10} \text{ cm}^{-3}$, corresponding to $f/i = 2.41$. Fig. 5.6 shows the fit for the low-density limit (solid histogram), and the 90% upper limit (dotted histogram). The strong $f$ line clearly requires a low-density environment; the 90% limit, while formally acceptable, requires $i > f$, unlike the data. The densities suggested here, $n_e < 10^{11} \text{ cm}^{-3}$, are typical of stellar coronae (e.g., Ness et al. 2001).

We now consider the influence of the radiation field. UV radiation will lower the $f/i$ ratio and thus simulate higher densities, i.e., the electron densities reported above are overestimated (Eq. 5.2). We now assume that the $f/i$ ratio is not suppressed by high electron densities and ask how far the source must be from the star in order to show the observed $f/i$ ratio. For $T_{\text{eff}} = 10050 \text{ K}$ (Table 5.1), we find from Eq. (5.2)–(5.5), that the 68% upper limit ($f/i = 2.41$) is attained at $d = 4.6R$, whereas the 90% upper limit ($f/i = 0.95$) requires
Chapter 5. The First High-Resolution X-Ray Spectrum of a Herbig Star: AB Aurigae

5.6 Discussion

In Table 5.1 we have listed the basic properties of AB Aur and we have compared them with those of the CTTS SU Aur and HD 163296, a Herbig star that has been reported to have a soft spectrum (Swartz et al. 2005).

$d > 2.1 R$. For $T_{\text{eff}} = 9500$ K reported by van den Ancker et al. (1998), the 68% upper limit corresponds to $d = 3.6 R$ and the 90% upper limit to $d = 1.7 R$. For a given $\mathcal{R}$ ratio, the origin of most of the emission must fulfill these requirements, but some contributions from closer to the star are not excluded.

We have assumed here that the radiation temperature is identical to $T_{\text{eff}}$ which, due to spectral modifications, may be inaccurate. Ness et al. (2001) found, for the relevant radiation at 1630 A, $T_{\text{rad}}$ several hundred K below $T_{\text{eff}}$ for nearby F and G-type stars, while Ness et al. (2002) reported $T_{\text{rad}}$ about 300 K below $T_{\text{eff}}$ for the B8 star Algol ($T_{\text{eff}} = 13000$ K). But even if we adopt $T_{\text{rad}} = 9000$ K for AB Aur, we still require $d > 2.8 R$ and $d > 1.3 R$ for the 68% and the 90% limit of the $f/i$ ratio.

Thus, we find that the electron density in the source must not exceed a few times $10^{10}$ cm$^{-3}$ (about $10^{11}$ cm$^{-3}$ for 90% upper limit) and the majority of the source plasma must not be closer to the stellar center than $(1.3 - 2.1) R$, depending on the exact temperature of the radiation field.
5.6. Discussion

The similarity between AB Aur and HD 163296 is striking. The stars not only have very similar fundamental properties, but their X-ray properties are also very similar. In contrast, SU Aur, a relatively massive but markedly later-type classical T Tau star, shows X-ray properties very different from the Herbig stars, although these properties are typical of T Tau stars (Güdel et al. 2007a).

We now discuss various models for X-ray production proposed in the literature and compare predictions made by these models with our spectroscopic results. We follow the presentation by Skinner & Yamauchi (1996).

5.6.1 Atmospheric and Wind Structure

Numerous optical and ultraviolet spectral observations of AB Aur have converged to a wind+chromosphere model in which an expanding wind overlies an extended, hot chromospheric layer. The latter has a height of about $1.5R$, with a temperature peak of $1.7 \times 10^4$ K (Böhm & Catala 1993). Praderie et al. (1986) and Catala et al. (1986) discovered periodic modulation in the Mg II and O II lines, although the periods disagreed, the period of Mg II being 45 hr and the period of O II being near 32 hr. The latter was identified with the stellar rotation period, implying formation of O II close to the star, while Mg II forms further away from the star (in its wind at several $R$) and may therefore be modulated by the rotation period of the envelope at that distance. The modulation was suggested to be due to a non-axisymmetric wind in which fast and slow streams alternate. Magnetic fields would then provide a possible explanation for this wind structure. High-temperature lines of N v and O vi were detected by FUSE, in AB Aur (Bouret et al. 1997; Roberge et al. 2001) and the similar HD 163296 (Deleuil et al. 2005), and these were interpreted as originating from shocks formed when the fast and slow winds collide (Bouret et al. 1997). The same model could also produce X-rays in a layer close to the star ($0.05R$ above the surface, Bouret et al. 1997). However, the $f/i$ ratio that we measured in the X-ray spectrum suggests that the X-rays are formed at distances $d > 1.3 R$ from the center of the star (for $T_{\text{rad}} > 9000$ K). Alternatively, N v and O vi could also be produced in a wind shock along with X-rays in the model by Babel & Montmerle (1997) described below.

Catala et al. (1999) extended periodicity studies to photospheric lines and found that the amplitudes of the red emission components are modulated with a period of 43 hr, whereas the velocity of the blue absorption components is subject to a 34 hr period. Further, the HeI D3 line shows red and blue components that are both modulated with a period around 45 hr (red in amplitude, blue in velocity). These authors interpreted the blue HeI D3 components as originating from an equatorial wind, whereas the photospheric blue absorption would come from high-latitude photospheric regions with radial flows, indicating a shorter period at high latitudes than near the equator. Finally,
all redshifted components would be due to polar infall. Their 43–45 hr period remains unexplained, but the infall may be magnetically linked to outflows at lower latitudes. Alternatively, the 45 hr HeI (blueshifted) outflow signatures may originate from a magnetic disk wind, indicating an anchor point of the magnetic fields at 1.6\(R\) (Catala et al. 1999).

The above interpretation is subject to one caveat, namely the assumed high inclination angle (70° in Catala et al. 1999). Adopting a radius of 2.3 \(R_\odot\) (Table 5.1), \(v \sin i = 80\) km s\(^{-1}\) (Böhm & Catala 1993, Table 5.1), and the extreme value of \(i = 70°\) (see also references in Corder et al. 2005), we find \(P = 33\) hr, in agreement with the 32-34 hr period reported from O\(\Pi\) and blue photospheric absorption (for \(i = 90°\), the maximum rotation period is 35 hr). However, recent observations have strongly revised the disk inclination angle and now suggest \(i \approx 21.5°\) (Corder et al. 2005 and further references therein). This suggests \(P = 12.9\) hr. The 32-34 hr stellar rotation period could only be maintained if the disk and stellar axis were grossly mis-aligned.

Our X-ray light curve period of 42.5 \(\pm\) 4 hr perfectly agrees with the period in Mg\(\Pi\) and He I, i.e., components formed in the chromosphere at the wind base and in the wind itself, but is clearly not compatible with the rotation period based on \(v \sin i\) measurements with \(i \approx 21°\), or the photospheric blue absorption components. We thus tentatively conclude that the X-ray production may be related, in some way, to the wind.

### 5.6.2 Wind Shocks

In hot stars, shocks are driven by instabilities in line-driven winds (see, for example, Feldmeier et al. 1997a for a review). The electron temperature of AB Aur’s X-ray source is similar to those measured in O stars (Feldmeier et al. 1997a). The important parameter for shock instability is the Eddington parameter

\[
\Gamma = \frac{\sigma_e}{2.5 \times 10^8 M} \frac{L_*}{M} \tag{5.6}
\]

(Castor et al. 1975) where we set the electron mass scattering coefficient \(\sigma_e = 0.4\) cm\(^2\) g\(^{-1}\), and \(M\) is the stellar mass. We thus obtain \(\Gamma = 5.6 \times 10^{-4} \ll 1\). Under these circumstances, the wind cannot be radiation-driven, which makes the instability-shock hypothesis unlikely.

On the other hand, Zinnecker & Preibisch (1994) have suggested that the wind slams into dense molecular material in the ambient medium. Using an average wind velocity of Catala & Kunasz (1987), \(v_w = 225\) km s\(^{-1}\), and the wind mass loss rate reported by Skinner et al. (1993), \(\dot{M}_w = 1.1 \times 10^{-8} M_\odot\) yr\(^{-1}\) from radio observations, we derive a “kinetic luminosity” \(\dot{M}_w v_w^2 / 2 \approx 1.8 \times 10^{32}\) erg s\(^{-1}\). That means that only about 0.3% of the available wind energy needs to be dissipated in shocks to produce the X-ray luminosity.
However, the observed systematic variability of the X-ray source on time scales of several hours makes models based on very-large-scale shocks as well as on many distributed shocks in the wind of a single star unlikely. This echoes the conclusions by Skinner et al. (2005) who summarize observations of X-ray variability in O stars. Although Feldmeier et al. (1997b) proposed variability owing to colliding shells in the O-star wind, the time scales for such collisions would be shorter (of the order of 500 s).

5.6.3 Magnetically Confined Winds

Although strong magnetic fields are not expected on Herbig stars given their predominantly radiative interior, large-scale fossil magnetic fields from the stellar formation process could still be trapped in the star. There have been a number of investigations that studied the consequences of ionized winds trapped in large-scale stellar magnetic fields.

Havnes & Goertz (1984) presented a model of a global stellar magnetosphere which is fed by a wind from the stellar surface. They considered a magnetic field with a dipole configuration, where plasma is confined by closed magnetic field lines. Close to the stellar surface, where the gravitational forces exceed the centrifugal forces, the density is low, while it increase further out. The temperature also increases, possibly to coronal temperatures, at 5-10 stellar radii. The outward transport of plasma takes place by events where the magnetic lines are broken due to excessive density. The energy input in the corona from rotation is given by 

\[ E = \frac{1}{2} \dot{M} \Omega^2 R^2 (L_2^2 - L_1^2) \]

where \( \Omega \) is the angular velocity. This can be rewritten as 

\[ E \approx 1.7 \times 10^{17} \dot{M} R^2 (L_2^2 - L_1^2) P_d^{-2} \text{ erg s}^{-1} \]

(\( \dot{M} \) is in \( M_\odot \) yr\(^{-1} \), \( R \) is in cm and the rotation period of the star \( P_d \) is in days). The parameters \( L_1 \) and \( L_2 \) are the distances to the inner and outer edges of the X-ray emitting region in units of the stellar radius. Both \( L_1 \) and \( L_2 \) are a function of the mass loss, the magnetic field and the rotation period, and are therefore difficult to estimate. Havnes & Goertz (1984) estimated these values to be \( L_1 = 15 \) and \( L_2 = 20 \), for a hotter and more massive star. Even if we cannot constrain these two values, we expect that \( (L_2^2 - L_1^2) \) will be larger than 1, so that we obtain \( E \geq 4 \times 10^{34} \text{ erg s}^{-1} \), i.e., enough to produce the observed \( L_\text{X} \).

In another model, Usov & Melrose (1992) considered the wind zone (open field lines) outside the corotating magnetosphere (dead-zone). According to their model a current sheet is formed in the equatorial plane outside the dead zone that separates regions with opposite directions of the magnetic field. The authors estimated the temperature and the energy released in the current sheet, assuming bremsstrahlung as a cooling agent. Assuming a surface magnetic field strength of 100 G (see recent measurements on Herbig stars by Hubrig et al. 2004), the above \( \dot{M} \), an average wind velocity of \( \dot{v}_w = 225 \text{ km s}^{-1} \) (Catala & Kunasz 1987; Skinner et al. 1993), and a radius of \( 2 - 4 R_\odot \), we find for the two
parameters $\xi$ and $\eta$ in their Eqs. (5) and (23), $\xi \gg 1$ and $\eta \geq 1$, respectively, and therefore, from Eq. (25) in Usov & Melrose (1992), an X-ray luminosity of $1.1 \times 10^{30}$ erg s$^{-1}$. Their Eq. (20) predicts an electron temperature of $T \approx 6 \times 10^7 (v_w/10^8 \text{ cm s}^{-1})^2 \approx 3 \times 10^6$ K. These parameters are again in good agreement with our measurements.

Finally, Babel & Montmerle (1997) considered the wind shock inside the magnetosphere that emerges when the magnetically guided winds from the northern and the southern hemispheres collide in the equatorial plane. They predict a shock temperature of $T_s = 1.13 \times 10^5 (v_w/10^7 \text{ cm s}^{-1})^2 \approx 0.5 - 1$ MK for $v_w = 200 - 300$ km s$^{-1}$ (extreme values reported by Catala & Kunasz 1987). Magnetically guided winds develop shocked equatorial “disks” only if the magnetic fields are sufficiently strong for confinement. For this, the (equatorial) wind confinement parameter, $\eta = B^2 R^2 / (\dot{M} v_W)$ must be at least unity (ud-Doula & Owocki 2002; note that $B$ is the surface magnetic field). With our stellar parameters (Table 5.1) and $B = 100$ G, we find $\eta \approx 20$, and hence wind shocks can develop. The temperatures provided by this model are nevertheless too low with respect to the observations.

In summary, wind-fed magnetospheres may be promising to produce the observed X-ray emission, although details of the magnetic field arrangement and the wind-field interactions would need further elaboration for the specific case of AB Aur.

### 5.6.4 Accretion-Induced X-Rays

Accretion has recently gained some attention as a possible contributor to X-ray emission in classical T Tau stars. Accretion shocks at the base of magnetic funnel flows may reach high temperatures, and high densities (of order $10^{13}$ cm$^{-3}$, Ulrich 1976). In standard accretion shock models, the shock heats up to a few times $10^6$ K, with the ensuing X-rays heating the underlying photosphere to produce an UV excess (Calvet & Gullbring 1998). Lamzin (1999) concluded that a typically small fraction of the X-rays escape from the shock that can be seen in soft X-rays. However, it is important to estimate whether the shock is above the photosphere at all and therefore visible, and this is not normally the case for T Tau stars with average accretion characteristics (Calvet & Gullbring 1998). Little variability should be seen in $L_X$ or the electron temperature in shocks (Lamzin 1999). Kastner et al. (2002) have proposed accretion-induced X-ray production for the unusually soft X-ray emission and the high densities measured in the spectrum of the classical T Tau star TW Hya. In analogy, Swartz et al. (2005) suggested the same scenario for the soft spectrum of the Herbig star HD 163296 (Table 5.1), but high-resolution X-ray spectroscopy was not available for this star.

A further argument in favor of an accretion model was brought up by Stelzer & Schmitt (2004) who argued that TW Hya’s anomalously high abundances
of Ne and N support an accretion scenario; depleted metals would condense onto grains in the disk, leaving gas enriched in certain other elements, and this gas would eventually accrete onto the star. However, referring to Ne/Fe and N/Fe abundance ratios, there are also evolved stars and non-accreting pre-main sequence stars that reveal high values for these ratios (see Güdel 2004a for a review). As for the Ne/O abundance ratio, Drake et al. (2005b) indeed found it to be unusually high in TW Hya compared to other stars. However, such a high Ne/O abundance ratio has so far been measured only in TW Hya. In other accreting and non-accreting stars, this ratio is found to be half as high as in TW Hya (see for example Argiroffi et al. 2005; Robrade & Schmitt 2006; and Chapter 3), and consistent with values of a large sample of late-type stars (Drake et al. 2005b; Chapter 2). Drake et al. (2005b) suggested that the anomalous Ne/O abundance in TW Hya is determined by the higher degree of metal depletion in this older star. This ratio can therefore not in general be used as an accretion signature in young and less evolved TTS.

What do we know about accretion in Herbig stars? Evidence for accretion from a disk is rather indirect. The temperature of the photosphere that is heated by infalling material is very similar to the undisturbed photosphere (Muzerolle et al. 2004). There has been some evidence for mass inflow in Herbig stars from redshifted absorption components in optical lines (Sorelli et al. 1996; Natta et al. 2000; Catala et al. 1999), but the interpretation is model-dependent, with \( \dot{M} \) possibly exceeding \( 10^{-7} M_\odot \text{ yr}^{-1} \) (Sorelli et al. 1996). Blondel et al. (1993) suggested that hydrogen Ly\( \alpha \) lines in Herbig stars are due to infalling gas, although no signatures were found for AB Aur. Grady et al. (1999) discussed evidence for infalling gas in the case of AB Aur. Muzerolle et al. (2004) interpreted Balmer and sodium profiles based on magnetospheric accretion models to conclude that \( \dot{M} \approx 10^{-8} M_\odot \text{ yr}^{-1} \) for the Herbig star UX Ori, and for a larger sample of stars that excess fluxes in the Balmer discontinuity imply \( \dot{M} \lesssim 10^{-7} M_\odot \text{ yr}^{-1} \). However, the Balmer discontinuity excess has been measured to be 0±0.1 mag in AB Aur (Garrison 1978), at best implying \( \dot{M} \approx 10^{-8} M_\odot \text{ yr}^{-1} \) (after Muzerolle et al. 2004). This is supported by measurements by Böhm & Catala (1993) who find \( M \lesssim 7.5 \times 10^{-8} M_\odot \text{ yr}^{-1} \). Catala et al. (1999) find explicit evidence for near-polar downflows in AB Aur with velocities of about 300 km s\(^{-1}\), leading to estimates for \( \dot{M} \) of a few times \( 10^{-9} M_\odot \text{ yr}^{-1} \) (see also Bouret & Catala 2000).

The accretion model is specifically favored for AB Aur by the measurements of the red component of the He I D3 line by Catala et al. (1999). They found a periodicity in the redshifted line amplitude of 42–45 hr, i.e. very close to the period that we measure in the X-ray light curve.

We first need to check whether an accretion rate of order \( \dot{M} \approx 10^{-8±1} M_\odot \text{ yr}^{-1} \) suffices to explain the observed X-ray output if the latter is indeed generated by accretion shocks. The accretion luminosity, assuming that the disk is truncated at the corotation radius (\( \approx 2R \) from the center of the
star for AB Aur), is \( L_{\text{acc}} = 0.5GM \dot{M}/R \approx 6 \times 10^{40} \dot{M} \dot{M}/(M_\odot \text{ yr}^{-1})/\dot{R} \text{ erg s}^{-1} \) or \( L_{\text{acc},30} \approx 600\dot{M}M_\odot/\dot{R} \) where \( L_{\text{acc},30} = L_{\text{acc}}/(10^{30} \text{ erg s}^{-1}) \), \( \dot{M} = M_\odot/M_\odot \), \( \dot{R} = R/R_\odot \), and \( \dot{M}_{8} \) is \( \dot{M} \) in units of \( 10^{-8} \) \( M_\odot \text{ yr}^{-1} \) (similar values were reported for HD 163296 given the very similar parameters for this star and AB Aur - see Table 5.1, Swartz et al. 2005). We conclude that there is sufficient accretion energy available to produce the X-rays.

We now estimate what an accretion model would predict for X-ray production on AB Aur. The expected shock temperature is, from the strong-shock conditions, \( T = 3v^2_m/m_p \) where the upstream flow velocity is approximately equal to the free-fall velocity, \( v = (2GM/R) \) (ignoring centrifugal forces if mass is guided along rotating magnetic fields), \( m_p \) is the proton mass, and \( \mu \approx 0.62 \) for a fully ionized gas. We thus find
\[
T \approx 5.4 \times 10^6 \frac{\dot{M}}{\dot{R}} \text{[K]} \tag{5.7}
\]
and with the parameters in Table 5.1, \( T \approx 6 \) MK, in good agreement with our measurements.

To estimate the shock density, we use the strong-shock condition \( n_2 = 4n_1 \) where \( n_1 \) and \( n_2 \) are the pre-shock and post-shock densities, respectively. We first estimate \( n_1 \) from the total mass accretion rate and the estimated accreting area on the surface: \( \dot{M} \approx 4\pi R^2 f v_\text{in} n_e m_p \) where \( f \) is the surface filling factor of the accretion flows, or \( \dot{M}_{8} \approx 1 \times 10^{-11} \dot{R}^{3/2} \dot{M}^{1/2} f n_1 \). We thus find\(^3\)
\[
n_2 \approx \frac{4 \times 10^{11} \dot{M}_{8}}{R^{3/2} \dot{M}^{1/2} f} \text{[cm}^{-3}] \tag{5.8}
\]
We adopt \( \dot{M}_{8} = 1 \), which seems to be a reasonable value based on the findings summarized above, and which is similar to accretion rates of less massive T Tau stars. For typical filling factors as discussed by Calvet \& Gullbring (1998), i.e., \( f = 0.1 - 10\% \), we find \( n_2 \approx 10^{12} - 10^{14} \text{ cm}^{-3} \).

Although these are densities similar to those measured on the T Tau star TW Hya, the O\textsc{vii} triplet we see in AB Aur requires densities about 100 times smaller. This could only be achieved by lowering \( \dot{M} \) to about \( 10^{-10} M_\odot \text{ yr}^{-1} \) or by increasing the accretion area to essentially the entire stellar surface. The former possibility is not supported by (at least tentative) measurements of limits to \( \dot{M} \), as summarized above. Accretion onto the entire stellar surface is unreasonable given that the star accretes from a disk, and a wind is present (e.g., Praderie et al. 1986).

\(^3\)In principle, \( n_2 \) could also be derived from the observed \( L_X \) and the shock volume, which derives from \( f \) and the shock height, the latter being dependent on the inflow velocity and the cooling time, see Calvet \& Gullbring (1998). However, such shock models show that the X-rays are attenuated by 1–2 orders of magnitude due to the falling material, see Lamzin (1999). Assuming that all \( L_X \) escapes would, when scaled with the observed luminosity, lead to an underestimated \( n_2 \) and an overestimated shock height.
We have not yet considered the radiation field of the A star. As shown in Sect. 5.5, a minimum distance of $1.3 - 2.1R$ from the center of the star is required for the X-ray source to be compatible with the observed O vii $f/i$ flux ratio. One way out is that the shocks are sufficiently shielded from UV radiation. It is not clear, then, how X-rays can escape without any absorption in addition to the circumstellar photoelectric absorption that agrees well with the optical extinction from circumstellar dust. This absorption is in fact extremely low in AB Aur, compared to other young stars in the Taurus-Auriga molecular cloud (Güdel et al. 2007a).

In summary, then, certain properties of X-ray production in accretion shocks may well be explained by simple shock models, but there are serious problems with this explanation, in particular related to i) the radiation field, ii) the low densities measured in the O vii triplet and iii) the lack of any excess absorption. There is also little support with regard to selective condensation of metals onto grains in the accretion disk. The abundance of Fe, thought to be among the elements that condense easily onto grains (Stelzer & Schmitt 2004), is higher in the X-ray source than in the photosphere (Acke & Waelkens 2004).

5.6.5 Coronal X-Rays

Coronae provide the standard explanation for X-rays from stars of spectral type F and later, because these stars maintain an outer convection zone that drives a dynamo. AB Aur is, however, a late-B or early-A type star.

Models of Herbig stars have indicated that a transient shell in which deuterium is burned may develop in Herbig stars, although this occurs preferentially in the later-type Herbig Ae stars (Palla & Stahler 1993). The calculations of Siess et al. (2000) of pre-main-sequence evolutionary tracks also predict the presence of a thin convective layer in young AB stars. Using their evolutionary model, we obtain for AB Aur a thin convection zone of 0.2% of the stellar radius.

The situation resembles that of mid-to-late A-type stars and early-F type stars on the main sequence. Although some of these stars are X-ray sources, they are clearly subluminous compared to later-type stars (if normalized with $L_\star$) despite their often rapid rotation, and the X-ray spectra are soft (Panzera et al. 1999). It appears that the thin convection zones of these stars are unable to maintain vigorous magnetic dynamos, resulting in soft, solar-like coronae. A similar situation may apply to Herbig stars like AB Aur.

The average temperature of AB Aur is similar to $T_\text{ac}$ of moderately active main-sequence solar analogs (G2-5 V), such as π1 UMa, χ1 Ori, and κ1 Cet (see Chapter 2). Assuming that AB Aur reveals a similarly structured corona, we expect that $L_X$ scales with the surface area. Adopting $R = 2.3R_\odot$ for AB Aur, its $L_X$ would be $\approx (5 - 6) \times 10^{29}$ erg s$^{-1}$, in agreement with the
observations. We also note that the densities derived from the O vii lines are in good agreement with measurements reported for numerous coronal sources (Ness et al. 2004). Finally, the modulation of the X-ray light curve observed in Fig. 5.1 could be due to rotation and is thus also consistent with the coronal hypothesis.

Tout & Pringle (1995) have proposed a non-solar dynamo that could operate in rapidly rotating A-type stars based on rotational shear energy. The model predicts, for the time development of the X-ray emission from the associated corona,

\[ L_X(t) = L_{X,0} \left(1 + \frac{t}{t_0}\right)^{-3} \]  

(5.9)

where \( L_{X,0} \) is a quantity that depends on stellar mass, radius, the amount of differential rotation in the stellar interior, the breakup velocity, the coronal heating efficiency, and the efficiency of magnetic field generation. Following the argumentation and the choice of constants given in Skinner et al. (2004) and Tout & Pringle (1995), we find \( L_{X,0} \approx 1.6 \times 10^{31} \text{ erg s}^{-1} \). For AB Aur, Eq. (3.15) in Tout & Pringle (1995) gives, using the same default parameters, \( t_0 = 6.3 \times 10^5 M^{-1/2} R^{3/2} \text{ yr} \approx 1.5 \times 10^6 \text{ yr} \). Given the star’s age of 4 Myr (DeWarf et al. 2003), we expect \( L_X \approx (3 - 4) \times 10^{29} \text{ erg s}^{-1} \). Using our updated parameters for AB Aur, we thus confirm the conclusion by Skinner et al. (2004) that this model provides very good agreement with the observations of AB Aur, although this is not true for most other Herbig stars. As the star evolves, the X-ray generation would rapidly decay further.

We found a coronal Fe abundance that is at least equal to the photospheric abundance. This is again consistent with a coronal model. In more evolved magnetically active stars, the Fe abundance generally increases toward lower activity and becomes comparable with the photospheric abundance in inactive stars (Gudel 2004a).

We note, however, that centrifugal forces exceed gravitational forces at a distance of 1.68 \( R \) from the center of the star according to parameters in Table 5.1 and assuming an (unconfirmed) rotation period of 12.9 hours. For equatorial magnetic fields, then, the coronal radius must be less than 1.68 \( R \), otherwise the loops will be unstable (Collier Cameron 1988). This condition is only in marginal agreement with the lower limit to the source size from the \( f/i \) ratio (Sect. 5.5) unless \( P_{\text{rot}} \) is larger.

### 5.6.6 X-Rays from a Companion?

A majority of stars form in multiple systems, and this is particularly true for Herbig stars (Feigelson et al. 2003). A companion may therefore also be responsible for the observed X-ray emission in AB Aur. Such a companion would most likely be a T Tau star. Recently, Stelzer et al. (2006) have studied
a sample of 17 Herbig Ae/Be stars with Chandra, concluding that at least in 7 stars the X-rays could originate from an unresolved companion. Only 6 stars are found to be X-ray emitters with no visual or spectroscopic detection. Furthermore, the X-ray properties in this stellar sample are very similar to X-ray properties of CTTS.

Behar et al. (2004) postulated that the X-ray emission detected from the late-type B star μ Lep in fact originates from an unknown pre-main-sequence companion, given the high $f/i$ ratio measured in its spectrum. For AB Aur, the $f/i$ ratio requires the bulk X-ray emission to originate at $r > 1.38R$. This would be fulfilled if the X-ray emission originated from a companion.

Stringent constraints have been discussed in the literature, interpreted and summarized by Piétu et al. (2005) (see their Sect. 5.1, and references therein): Any co-eval companion within 120-1500 AU (0.86-12.5") must have a mass $< 0.02M_\odot$. The mass upper limit is $0.25M_\odot$ down to a separation of 0.4". In the range of 0.07–10", the upper limit to a companion mass is, from speckle interferometry, 0.05–0.3$M_\odot$. Recently, Baines et al. (2006) reported evidence for binarity of several Herbig stars, including AB Aur. However, they estimated a separation of 0.5-3.0" for AB Aur.

The point-spread function of XMM-Newton does not allow us to distinguish between our target source and potential companions within a few arcsec. We have therefore analyzed an exposure of the region around SU Aur + AB Aur obtained from the archive of the Chandra X-Ray Observatory, revealing much better positional information (obs ID = 3755). Standard data reduction methods and up-to-date aspect corrections were applied. The observation used the ACIS-S detector, with the high-energy grating inserted. It was centered on SU Aur, with AB Aur being located close to the chip edge. Using the wavdetect (wavelet detection) routine in the Chandra CIAO software, we measured the centroid positions of the images of both stars, to find

<table>
<thead>
<tr>
<th>Star</th>
<th>RA(J2000.0)</th>
<th>δ(J2000.0)</th>
</tr>
</thead>
<tbody>
<tr>
<td>AB Aur</td>
<td>$4^h55^m45.846^s\pm0.17^s$</td>
<td>$30^\circ33'04.138''\pm0.17''$</td>
</tr>
<tr>
<td>SU Aur</td>
<td>$4^h55^m59.389^s\pm0.11^s$</td>
<td>$30^\circ33'04.1297''\pm0.11''$</td>
</tr>
</tbody>
</table>

The nearest 2MASS objects are located at

<table>
<thead>
<tr>
<th>Star</th>
<th>RA(J2000.0)</th>
<th>δ(J2000.0)</th>
</tr>
</thead>
<tbody>
<tr>
<td>AB Aur</td>
<td>$4^h55^m45.826^s\pm0.07^s$</td>
<td>$30^\circ33'04.37''\pm0.07''$</td>
</tr>
<tr>
<td>SU Aur</td>
<td>$4^h55^m59.381^s\pm0.06^s$</td>
<td>$30^\circ34'01.56''\pm0.06''$</td>
</tr>
</tbody>
</table>

The position offsets of the X-ray sources are thus 0.258" and 0.103" in RA for AB Aur and SU Aur, respectively, and $-0.232''$ and $-0.262''$ in declination. The similar offsets for both stars suggest a systematic offset of the pointing
of order 0.28″, well within the errors of the Chandra attitude solution.\footnote{The 90\% source location error circle in Chandra has a radius of about 0.5″, see Chandra Proposers’ Observatory Guide v.8} Correcting the AB Aur position by the offset of the better determined SU Aur position, the deviation from 2MASS of AB Aur is only 0.16″ in RA and 0.03″ in declination, which is within the errors of the measurement. We conclude that we have identified AB Aur in X-rays well within 0.5″ (at the 3σ level) of the 2MASS position. It is therefore improbable that the source of the X-rays is the companion detected by Baines et al. (2006), which is thought to have a separation with AB Aur of > 0.5″.

Further arguments favor intrinsic X-ray emission from the Herbig star. In particular, the observed low average temperature is rather uncommon to lower-mass T Tau stars; the latter rather show hot components with characteristic temperatures up to 20-30 MK as, for example, SU Aur but also other CTTS in the XEST survey (see previous Chapters). Very-low mass stars and brown dwarfs do reveal softer spectra with dominant temperatures below 10 MK but their total X-ray luminosities are much below $L_X$ measured here for AB Aur (see Grosso et al. 2007a for the brown dwarf sample from the XEST survey).

We also note that the $N_H$ value determined in the XEST survey ($N_H \approx 5 \times 10^{20}$ cm$^{-1}$, Gudel et al. 2007a) is in perfect agreement with the visual extinction measured for AB Aur by Roberge et al. (2001) (0.25 mag), if we apply a standard interstellar conversion law, $N_H \approx 2 \times 10^{21} A_V$ mag (Vuong et al. 2003 and references therein).

The most substantial argument against the companion hypothesis is the close coincidence between the X-ray period and the period observed in the lines of the wind of the Herbig star.

Finally, AB Aur and HD 163296 are very similar in both their intrinsic properties and their X-ray properties. If the X-rays indeed originate from nearby T Tauri companions, then the companion of AB Aur would happen to be very similar to the companion of HD 163296.

Taken together, these arguments suggest that the X-rays are not originating from a companion but from AB Aur itself. This is different from flaring, hard sources among Herbig stars for which T Tau companions have recently been identified (see Sect. 5.1). We therefore suggest that the unusually soft emission is indeed a distinguishing property of genuine Herbig star X-ray emission.

### 5.6.7 Disk Related Models

The X-ray luminosities of A and B stars decrease as they approach and reach the main sequence (Stelzer et al. 2006); at the same time the outflow activity is believed to cease and the dense surrounding gas dissipates. That suggests that the X-rays of AB Aur could be related to the presence of the circumstellar
disk. Two different models that link the X-ray emission with the presence of the disk have been discussed in the literature: a disk corona and reconnection of magnetic fields that link the star to the disk. We describe these models, although they do not currently make predictions that we can test against our data.

The presence of a disk corona was discussed by Zinnecker & Preibisch (1994), but very little is known about its generation. The ionization due to the decay of radioactive nuclides could increase the conductivity of the disk to a level high enough to generate a magnetic field. The differential rotation in the disk could then generate a disk corona.

X-ray emission generated by reconnection of magnetic fields linking the star to the disk has been discussed as a model for the X-ray generation mechanism of low-mass protostars (Montmerle et al. 2000). If the rotation period of the star is not the same as the rotation period of the disk, a magnetic loop connecting the disk with the star will twist and inflate until it comes into contact with itself and reconnects. A similar model could apply to Herbig stars, although we would expect higher, flare-like temperatures due to the reconnection process.

5.7 Conclusions

We have presented the first high-resolution X-ray spectrum of an Herbig Ae/Be star, namely AB Aur. The use of high-resolution spectroscopy has allowed us to obtain important spectral information that cannot be addressed with EPIC spectra alone. The O\textsuperscript{vii} triplet constrains the electron density and is therefore important for the discussion of the different models. Further, we have been able to reliably determine the abundances of the high-FIP elements O and N. Finally, the O\textsuperscript{viii}, O\textsuperscript{vii} and N\textsuperscript{vii} lines permitted us to constrain the cool plasma.

We found the X-ray spectrum of AB Aur to be rather soft, with spectral-fit results that are consistent with a mean coronal temperature of about 5 MK, i.e., much less than the usual temperatures of coronae of low-mass pre-main-sequence stars that usually exceed 10 MK. We found an X-ray luminosity of about $4 \times 10^{29}$ erg/s in the 0.3–10 keV range. We derived the abundances and found them not to follow a First Ionization Potential (FIP) distribution, nor an inverse FIP distribution. We normalized the coronal abundances to the new photospheric abundances of AB Aur found by Acke & Waelkens (2004), who measured a very low photospheric Fe abundance. The Fe coronal abundance then is at least as large as the photospheric value.

The density-sensitive O\textsuperscript{vii} triplet has been studied in detail. Although its S/N is moderate, we found that the line flux ratios indicate densities $n_e \lesssim 10^{11}$ cm$^{-3}$, with the best-fit value being at the low-density limit ($n_e \approx 10^{10}$ cm$^{-3}$), similar to what is commonly found in stellar coronae (Ness et al. 2004).
We have discussed several X-ray generation mechanisms, and provided supporting evidence or pointed at problematic features for each of them. First, the probability that the X-ray emission originates from a companion TTS is small. The X-ray source is identified within 0.5″ of the 2MASS source corresponding to AB Aur in the Chandra image, so that the companion would have a mass < 0.3$M_\odot$ according to the constraints given by Piétu et al. (2005). Such a low-mass star would rarely produce an X-ray luminosity as high as observed for AB Aur. Furthermore, the close coincidence of the period that we have measured in X-rays with the period measured in lines formed in the wind of AB Aur itself makes the companion hypothesis improbable.

Accretion-induced X-ray emission has been widely discussed in the literature. With the observed electron density, this would be possible only if $M \approx 10^{-10} M_\odot$ yr$^{-1}$, i.e. lower than the value suggested in the literature ($M \approx 10^{-8 \pm 1} M_\odot$ yr$^{-1}$), and a filling factor $\gtrsim 10\%$. The major problem with this scenario is that the UV radiation field (of the stellar photosphere and the shock itself) would suppress the forbidden line of the O$^{\text{vii}}$ triplet.

The hypothesis that X-ray emission is generated by shocks in a line-driven wind, similarly to the mechanism that is believed to produce X-rays in O stars, is ruled out by the observed variability and the inability of the radiation field to drive the wind.

Magnetic fields have recently been detected on several Herbig stars (Donaì et al. 1997; Hubrig et al. 2004; Wade et al. 2005). Two further possible mechanisms are fundamentally dependent on the existence of a magnetic field: coronal emission and magnetically confined winds.

Coronal emission of the type seen in the Sun requires the presence of a dynamo. The recent calculations by Siess et al. (2000) predict a thin convective layer for Herbig stars, which in the case of AB Aur amounts to 0.2% of the radius. The question then is whether this thin convection layer would be sufficient to generate the dynamo that results in a corona with the observed X-ray properties. The corona should be quite extended in order to allow the $f/i$ flux ratio in the O$^{\text{vii}}$ triplet to be larger than unity. This is plausible because the surface gravity of AB Aur is about half that of the Sun.

Magnetically confined winds are a promising alternative, the more so as winds have explicitly been measured (Catala et al. 1999). There is evidence supporting a model of this kind. The period of the modulation that we measured in the X-ray light curve (42.2 hr, Sect. 5.4.1) is very close to a modulation period of Mg$\text{II}$ lines thought to form in the wind (Praderie et al. 1986). Whatever the production mechanism of the X-rays, it is possible to assume that they are closely related to the wind or are produced by it. The advantage of such a model is that X-rays are produced at some distance from the stellar surface, which alleviates the problem of the suppression of the O$^{\text{vii}} f$ line flux.

In both these models, we encountered a problem when we adopted a stellar inclination angle of $i \approx 21.5^\circ$ (Corder et al. 2005) and $v\sin i \approx 80$ km s$^{-1}$
5.7. Conclusions

(Böhm & Catala 1993). The rotation period would then be 12.9 hr. This value is significantly smaller than both the MgII and X-ray period (≈ 42 hr) and the value suggested by Catala et al. (1999) for the rotation period (32-34 hr). Further, in the hypothesis of a stellar corona, the latter should in this case not exceed a height of 0.68 \( R \) (the location of the co-rotation radius where centrifugal forces \( \Omega d^2 \) equal gravitational forces \( GM/d^2 \)) above the surface at the equator, because otherwise centrifugal forces would make the loops unstable (Collier Cameron 1988). Further, with the adopted small inclination angle, no significant rotational modulation should occur, and the \( f/i \) ratio is at risk of being significantly suppressed.

The problems with a wind in a magnetosphere or a corona would be alleviated if the stellar rotation axis were, for some reason, inclined against the axis of the circumstellar disk, or the latter were largely warped between the inner regions and the well-observed outer regions (Corder et al. 2005). The 33 hr period suggested by Catala et al. (1999) to be identified with the rotation period would then require a stellar inclination of about 70 deg, making partial eclipses of X-ray emitting material, located for example close to the equatorial plane of the star, easily possible. At the same time, azimuthally varying wind velocities would produce the line shift periodicity as reported in the optical and UV (Praderie et al. 1986; Catala et al. 1986). However, we have no explanation as to why the rotation axis should be inclined against the disk.

We further note that HD 163296 displays very similar properties to AB Aur (Table 5.1), with a soft X-ray spectrum similar to the one that we have found for AB Aur. However, the X-ray properties found for AB Aur and HD 163296 are not common to all Herbig stars. Some of these stars display harder spectra, with temperatures reaching several tens of MK and with larger X-ray luminosities (Hamaguchi et al. 2005; Skinner et al. 2004; Stelzer et al. 2006). Given the peculiar properties of the X-rays of AB Aur and the similar HD 163296 (Table 5.1, Swartz et al. 2005), namely a very soft X-ray spectrum with a moderate \( L_X \), we suggest that these properties are defining properties of X-ray emission intrinsic to Herbig stars, while hard, flaring emission may be due to undetected companions (see Sect. 5.1, and Sect. 5.6.6).
Chapter 5. The First High-Resolution X-Ray Spectrum of a Herbig Star: AB Aurigae
Chapter 6

XMM-Newton Survey of the Chamaeleon I Star Forming Region

ABSTRACT: We present initial results of an XMM-Newton survey of the Chamaeleon I star forming region. In the five exposures we detected a total of 449 X-ray sources. Among them, 96 are counterparts to the 138 spectroscopically confirmed members in the surveyed region; six of them have a spectral type later than M6.5 and are therefore classified as brown dwarfs. The spectra of the brighter sources were analyzed using a continuous emission-measure model composed of two power laws in temperature, as suggested from high-resolution X-ray spectroscopy. We investigate correlations between the derived X-ray luminosity and stellar properties such as rotation rate, mass and bolometric luminosity. We find a deviation from the galactic $N_H/A_J$ relation in Chamaeleon I that could be explained with a larger characteristic dust-grain size.
6.1 Introduction

X-ray surveys play a crucial role in the study of high-energy radiation processes relevant to star and planet formation. Young stars, independently of their accretion rate, are very strong X-ray emitters.

The Chamaeleon I cloud is one of the nearest and best studied star forming regions at a distance of about 160 pc (Whittet et al. 1997). Its high galactic latitude ($b \approx -16^\circ$) minimizes the number of foreground and background stars, and its compactness allows the cloud to be surveyed with only a few XMM-Newton observations.

Our survey is based on five XMM-Newton observations that are shown in Figure 6.1 and summarized in Table 6.1. Using the maximum likelihood detection algorithm eboxdetect and emldetect within the XMM Newton SAS software, a total of 449 X-ray sources were detected in the five fields.

Figure 6.1: ROSAT contour map of the Chamaeleon I star forming region adapted from Feigelson at al. (1993) with the five XMM-Newton EPIC fields overplotted
Table 6.1: Observation log for the 5 Chamaeleon I fields

<table>
<thead>
<tr>
<th>Field</th>
<th>Pointing RA</th>
<th>Pointing dec</th>
<th>Exposure [sec]</th>
</tr>
</thead>
<tbody>
<tr>
<td>A</td>
<td>10:59:07.10</td>
<td>-77:01:49.0</td>
<td>35423</td>
</tr>
<tr>
<td>B</td>
<td>11:07:52.90</td>
<td>-77:36:56.0</td>
<td>31122</td>
</tr>
<tr>
<td>D</td>
<td>11:11:30.00</td>
<td>-76:31:00.0</td>
<td>26035</td>
</tr>
<tr>
<td>G</td>
<td>11:00:06.00</td>
<td>-77:28:51.0</td>
<td>29034</td>
</tr>
<tr>
<td>H</td>
<td>11:06:05.00</td>
<td>-77:10:40.0</td>
<td>28834</td>
</tr>
</tbody>
</table>

Using the works of Luhman 2004 (optical spectroscopy), Comeron et al. 2004 (Hα emission and spectroscopy) and Gomez & Mardones 2003 (NIR spectroscopy), we cross-identified our X-ray source catalog with the 138 known Chamaeleon I members in our observation fields. Of the X-ray sources, 96 are identified with Chamaeleon I members.

6.2 Model for Spectral Fit

Low-resolution spectra are often fitted using a 1-, 2- or 3-temperature model. From high-resolution spectroscopy, however, it has become evident that coronae display continuous emission measure distributions (EMD). The derivation of a continuous EMD is more ambiguous from low-resolution spectra such as those from the EPIC spectra. From a physical point of view, coronal, magnetically trapped plasmas are also expected to be arranged in continuous EMDs. We therefore use an EMD model as described in Chapter 1 (Sect. 1.6). According to this model, the EMD is composed of two power laws in temperature and is given by the formula:

\[
Q(T) = \begin{cases} 
EM_0 \cdot (T/T_0)^\alpha & \text{for } T \leq T_0 \\
EM_0 \cdot (T/T_0)^\beta & \text{for } T > T_0 
\end{cases}
\]  

(6.1)

where \(T_0\) is the temperature at the EMD peak and \(EM_0\) is its emission measure. The slope of the power law below \(T_0\) is \(\alpha\), whereas \(\beta\) is the power-law slope above \(T_0\). We introduce a low-temperature and an high-temperature cutoff, fixed at \(\log T = 6.0\) and \(\log T = 8.0\), respectively. We fixed \(\alpha\) at a value of 2, consistent with EMD slopes for young stars as found from high-resolution spectroscopy (see for example Argiroffi et al. 2004 and Chapter 2). Guided by the coronal element abundances found in the literature (Scelsi et al. 2005, Argiroffi et al. 2004, García-Alvarez et al. 2005 and Chapter 2) we adopted an abundance pattern typical for young stars (see Sect. 1.6). We fitted all,
except some very faint, sources in our five fields using this model, obtaining a homogeneous set of X-ray properties that can be used for studies of correlation with other stellar parameters.

6.3 Results

6.3.1 The Rotation-Activity Relation

In Figure 6.2 we plot $L_X/L_{\text{bol}}$ as a function of the rotation period $P_{\text{rot}}$. We divide our stellar population in two subpopulations, the classical T Tauri Stars (CTTS) that show evidence for strong accretion such as broad Hα emission lines, and the weak-lined T Tauri Stars (WTTS), that show only very weak Hα emission. We used the Hα equivalent widths from Luhman (2004) and references therein. Big and small circles represent CTTS and WTTS, respectively. Filled circles are bright sources with more than 100 counts, whereas empty circles are sources with less than 100 counts in the fitted spectrum. The asterisks mark the two Brown Dwarfs (BD) for which the rotation period is known. On the left plot the arrows indicate upper limits to the rotation periods that we derived from the $v\sin i$ values. The grey bar represents an average rotation-activity relation for main-sequence stars. This relation was empirically derived from the work of Pizzolato et al. (2003): the locus for the kink was calculated for K7-M0 main-sequence stars, equivalent to stars with a bolometric luminosity of $\approx 0.1L_\odot$. In Chamaeleon I we do not find a trend of decreasing $L_X/L_{\text{bol}}$ for slower rotators. On the contrary, as shown in the right panel of Figure 6.2, if we fit our data with a regression line, we find that $L_X/L_{\text{bol}}$ increases for higher
6.3. Results

The X-ray luminosity as a function of the bolometric luminosity. Key to the symbols: Size, from largest to smallest: CTTS - WTTS - unknown class. Asterisks are BDs. Asterisks with arrows are upper limits for undetected BDs. The open circles are faint sources, with less than 100 cts in the spectrum.

rotation periods. The regression line has a slope of 0.8 and is described by the relation $\log(L_X/L_{bol}) = -3.77(\pm 0.26) + 0.80(\pm 0.15) \cdot \log P_{rot}$. A similar trend was found in Orion by Preibisch et al. (2005): they reported a slope of 1.27 and a linear regression $\log(L_X/L_{bol}) = -4.21(\pm 0.07) + 1.27(\pm 0.09) \cdot \log P_{rot}$. In contrast, Güdel et al. (2006b) find for the Taurus Molecular Cloud decreasing $L_X/L_{bol}$ for longer rotation periods.

Unfortunately, in Cha I, the number of known rotation periods is small and the statistics is not sufficient to draw firm conclusions about the activity-rotation relation. We can state, however, that no appreciable drop of $L_X/L_{bol}$ is visible in the range of $P_{rot} = 5 - 10$ days.

6.3.2 The $L_X/L_{bol}$ Relation

In Figure 6.3 the X-ray luminosity is plotted as a function of the bolometric luminosity. Overplotted are the three lines for $L_X/L_{bol} = 10^{-4}$, $L_X/L_{bol} = 10^{-3}$, and $L_X/L_{bol} = 10^{-2}$. We note that nearly all low mass stars ($M < 2M_\odot$) have $L_X/L_{bol} > 10^{-4}$. Although the detected brown dwarfs (asterisks) are higher in $L_X/L_{bol}$ (mean value above $10^{-3}$) than low-mass stars (mean value of $10^{-3.3}$), the upper limits of undetected brown dwarfs fill the space below the detected ones. Thus we conclude that brown dwarfs are likely to follow the same trend as low mass stars in Chamaeleon I.
In order to estimate the mean value of $L_X/L_{bol}$ for the two populations of CTTS and WTTS, we plot histograms for $\log L_X/L_{bol}$ and fit the resulting distribution with a Gaussian. As shown in Figure 6.4, we find a mean value for $\log L_X/L_{bol}$ of approximately -3.3 for both CTTS and WTTS. We conclude that there is no global influence of the accretion behavior on the X-ray output of T Tau stars in ChaI.

### 6.3.3 The Mass-Activity Relation

In Figure 6.5 the X-ray luminosity is plotted as a function of the logarithm of the stellar mass. The masses are derived from the effective temperature and the bolometric luminosity using the isochrones of Siess et al. (2000) for masses larger than 0.1 $M_\odot$ and from Chabrier et al. (2000) for smaller masses. It is evident that $\log L_X$ increases with mass. The dash-dotted line describes the linear regression fit obtained by considering the stars between 0.1 and 2 $M_\odot$ (i.e. excluding the brown dwarfs). This regression line has a slope of 1.31 and is described by $\log L_X = 30.32(\pm0.1) + 1.31(\pm0.2) \cdot \log(M/M_\odot)$. We notice that the relation does not seem to be linear but is flatter for stars with higher masses. Again the BDs (detected and upper limits) fit well in the stellar relation. An X-ray luminosity-mass correlation was already found by Feigelson at al. (1993) in Chamaeleon I. In that study, based on ROSAT observations, a much steeper slope of 3.6 was found for the same correlation. Our dependence is, however, similar to the one found for the Orion Nebula Cluster by Preibisch
6.3. Results

et al. (2005), where a slope of 1.44 was reported.

6.3.4 The $N_H/A_J$ Relation

We plot in Figure 6.6 the hydrogen column density $N_H$ as a function of the infrared extinction $A_J$. In order to avoid effects from circumstellar material we concentrate in this study only on WTTS. The two thin lines are theoretical values for the $N_H/A_J$ relation valid for the interstellar medium, and they are calculated as follows. The ratio $N_H/A_V$ is taken from the literature: $N_H/A_V = 1.8 - 2.2 \times 10^{21}\text{cm}^{-2}/\text{mag}$. We then used the Cardelli et al. (1989) extinction law, for which the $A_J/A_V$ relation depends only on one parameter, $R_V = A_V/(A_B - A_V)$, and is described by the formula

$$A_J/A_V = 0.4008 - 0.3679/R_V.$$  

The galactic value for $R_V$ is found to be 3.1. Using this relation, we obtain

$$(N_H/A_J)_{\text{gal}} = 6.4 - 7.8 \times 10^{21}\text{cm}^{-2}/\text{mag}. \quad (6.3)$$

This $N_H/A_J$ range is plotted as a black line in both Figures 6.6. We see that in Cha I many stars lie on a line that indicates reduced $N_H$. This line is represented by the thick line that fits the relation

$$N_H/A_J = 4.1 \times 10^{21}\text{cm}^{-2}/\text{mag}. \quad (6.4)$$
A deviation from the galactic value of $N_H/A_J$ was also found and studied in detail by Vuong et al. (2003) for the star-forming region around ρ Oph. They obtained $N_H/A_J = 5.57 \times 10^{21}$ cm$^{-2}$/mag, using the Cardelli et al. (1989) $A_J/A_V$ relation. In their work, the measured $N_H/A_J$ value was used in a physical model of gas and dust properties to constrain the grain size distribution and the gas-to-dust ratio. They found a bigger mean grain size ($<a> = 0.035 - 0.095$ μm instead of $<a> = 0.008$ μm) and a gas-to-dust mass ratio of 80-95.

In Chamaeleon I, this model cannot be used, because the $N_H/A_J$ ratio is too small to derive a reasonable value for $R_V$ using the Cardelli et al. (1989) extinction curves (note that higher $R_V$ values were already claimed in Cha I, see for example Whittet et al. 1997 and Luhman 2004). However, we can state that this deviation from $(N_H/A_J)_{\text{gal}}$ is in the same direction of the deviation found by Vuong et al. (2003) and would also predict a higher $R_V$ value and a larger characteristic mean grain size. As the gas-to-dust ratio strongly depends on both $R_V$ and $N_H/A_J$ in the model, we are not able to predict the variation of gas-to-dust mass ratio in our case.

6.3.5 Summary and Conclusions

In the five XMM-Newton observations of the Chamaeleon I region we detected 96 cloud members. Of them, 6 are considered brown dwarfs. We correlated the X-ray activity, i.e. the ratio between X-ray luminosity and bolometric luminosity, with the rotation period, but we could not find any decrease in activity for slower rotators, as was found for main sequence stars. This result could be evidence for long convective turnover times in our sample of stars. We further studied the relation between X-ray luminosity and bolometric lu-
minosity and found that almost all low-mass stars have a ratio \( \log L_X / \log L_{\text{bol}} \) between \( -4 \) and \( -2 \), with a mean value at \( -3.3 \) for both CTTS and WTTS. We concluded that there is no influence of accretion on the X-ray output. An evident increase of \( L_X \) with the stellar masses is also found. Brown dwarfs do not show exceptional properties, but fit well to all the studied correlations. Finally, we investigated the ratio between the hydrogen column density \( N_H \) and the infrared extinction \( A_J \). We found \( N_H / A_J = 4.1 \times 10^{21} \text{cm}^{-2} / \text{mag} \), much lower than the galactic value \( N_H / A_J = 6.4 - 7.8 \times 10^{21} \text{cm}^{-2} / \text{mag} \) valid for the interstellar medium. This deviation is too large to draw quantitative conclusions about the gas-to-dust ratio, but we can qualitatively affirm that deviations in the same direction could be modeled with a larger characteristic mean grain size of the dust.
Chapter 7

Summary and Conclusions

In this thesis we have studied the coronal properties, inferred from X-ray spectra, of a sample of solar-like stars at different evolutionary stages. Spectra of main-sequence stars with different ages and spectra of pre-main sequence stars have been analyzed with the aim of reconstructing the history of the X-ray output of the Sun in time.

In Chapter 2 we analyzed six main-sequence solar analog stars with different ages. The youngest stars, 47 Cas B and EK Dra, have ages of 0.1 Gyr and have just arrived on the main sequence, while the oldest star, β Com has an age of 1.6 Gyr. As the stars age, their rotation rates decrease and their internal dynamos weaken, resulting in a decrease of the magnetic activity. As a consequence we found the stellar luminosity to decrease from a level very close to saturation in the zero-age main-sequence stars (ZAMS, \( L_X/L_{bol} \approx 10^{-3} \)) to a level of \( L_X/L_{bol} \approx 10^{-5} \) within the first 2 Gyr. During the next \( \approx 3 \) Gyr, as the solar age is approached, the X-ray luminosity decreases by another factor of 10. Evolutionary trends have also been found in the emission measure distributions: as a star ages, the temperature of the EMD peak decreases in time (from \( \approx 10 \) MK for 47 Cas B and EK Dra to \( \approx 4 \) MK for β Com). We thus found a power-law dependence between the average coronal temperature and the X-ray luminosity, namely, \( T_{av} \propto L_X^{0.25\pm0.02} \). This trend cannot be explained by a simple model in which the filling factor of the surface magnetic field increases in more active stars, as the larger filling factor would not explain the higher temperatures. Alternatively, an extreme model in which the coronal emission is entirely due to flaring loops was discussed. In this model, more active stars display a higher flare rate. However, the reason for this behavior is not fully clear. One possibility is that more active stars have a larger magnetic surface filling factor, so that the higher density of magnetic loops leads to more magnetic reconnection episodes, producing a higher rate of large flares (Güdel et al. 1997a).

Evolutionary effects were also found in the coronal abundances: as the
stars age, the abundances change from a pattern where abundances with a high First Ionization Potential (FIP) are enhanced with respect to elements with a low FIP (inverse-FIP effect) to a pattern analogous to the one observed in the Sun, where the low-FIP elements are enhanced relative to the high-FIP elements (FIP effect). This transition occurs at ages less than 300 Myr.

In Chapters 3-5, we studied the X-ray emission of pre-main sequence stars in the Taurus Molecular Cloud. Our studies were mainly concerned with T Tauri stars, and especially with the influence of accretion on the X-ray output. Using medium-resolution CCD spectra, we found that both accreting (classical T Tauri stars, CTTS) and non-accreting (weak-line T Tauri stars, WTTS) stars are saturated in X-rays, but saturation levels are different, with $L_X/L_{bol} = 10^{-3.73\pm0.05}$ for CTTS and $L_X/L_{bol} = 10^{-3.39\pm0.06}$ for WTTS. Given the similar distribution of $L_{bol}$ of the two stellar samples, the X-ray luminosity of CTTS is reduced by a factor of $\approx 2$ on average with respect to the X-ray luminosity of WTTS. Further, for WTTS a power-law correlation was found to apply between the average coronal temperature and $L_X$, namely, $T_{av} \propto L_X^{0.23\pm0.03}$, in agreement with the relation found in main-sequence stars. Such a correlation was found not to be present in CTTS; rather, CTTS coronae are generally hotter than those of WTTS. The lower X-ray production in CTTS refers to the range of plasma temperatures accessible with CCD spectra. It is therefore possible that part of the energy release in CTTS is shifted to lower temperature and therefore becomes inaccessible to the CCD detectors. Moreover, photoelectric absorption is typically larger in CTTS than in WTTS by a factor of $\approx 2.5$ due to the larger amount of circumstellar material, which makes the detection of soft emission more difficult.

The hypothesis that part of the emission is shifted to lower temperatures is in agreement with results we obtained from studying a sample of pre-main sequence stars with high-resolution spectroscopy. XMM-Newton RGS spectra are more sensitive to cool plasma and, using the flux ratio between the O vii lines and the O viii Ly$\alpha$ line, we discovered a soft excess in the spectra of accreting stars that is not present in the spectra of WTTS (Chapter 3). Possibly, this soft excess could account for at least some of the missing $L_X$ in CTTS although this could not be confirmed for two of the studied objects.

The cause of the additional soft component remains to be identified. One possible explanation is that accreting matter penetrates into the hot coronal magnetic structures and thus cools the plasma because more material needs to be heated, and because higher densities make the cooling more efficient. This process would cause a soft excess or would cool the plasma even further so that it is no longer accessible by X-ray detectors; it could therefore result in a deficiency of X-ray emission from accreting stars.

Previous studies of high-resolution spectra of CTTS have suggested the presence of X-ray emission formed at the base of the accretion stream. An accretion shock would also produce soft plasma that again could remain unde-
ected in CCD spectra. Further, an accretion shock would also result in higher electron densities. High electron densities have been found in the Chandra and XMM-Newton grating spectra of TW Hya ($n_e \approx 10^{13}$ cm$^{-3}$, Kastner et al. 2002; Stelzer & Schmitt 2004), BP Tau ($n_e \approx 10^{11}$ cm$^{-3}$, Schmitt et al. 2005; Robrade & Schmitt 2006, Chapter 3), and V4046 Sgr ($n_e \approx 3 \times 10^{13}$ cm$^{-3}$, Günther et al. 2006). On the other hand, the density-sensitive O VII triplet in T Tau is consistent with low densities ($n_e \approx 10^{10}$ cm$^{-3}$, Güdel et al. 2007b).

Similarly, we have found the spectrum of the accreting star AB Aur to be consistent with low densities (see Chapter 5). However, AB Aur is an Herbig Ae/Be star, and the mechanism for the generation of X-rays in these objects is poorly understood (Chapter 5). In conclusion, the role of accretion shocks in the X-ray output of CTTS remains controversial: while it seems to be a viable mechanism to produce the soft excess in some stars, it is not consistent with observations of all accreting stars.

Abundance anomalies have been found in pre-main sequence stars as well. A very high Ne/Fe abundance ratio was previously reported in the spectrum of TW Hya (Ne/Fe $\approx 7 - 11$, Kastner et al. 2002; Stelzer & Schmitt 2004), and interpreted as being due to metal depletion onto grains in the accretion disk. However, quite high Ne/Fe ratios have been measured in other pre-main sequence stars, like TWA 5 ($\approx 6$, Argiroffi et al. 2005), HD 98800 ($\approx 4$, Kastner et al. 2004), and BP Tau ($\approx 5$, Robrade & Schmitt 2006), where HD 98800 is believed to be a WTTTS. In Chapter 3 we have found high Ne/Fe abundance ratios (Ne/Fe $\approx 4 - 6$) for all stars except those with spectral type G (SU Aur, HP Tau/G2, and HD 283572). We therefore suggest that a high Ne/Fe abundance ratio is not a characteristic solely of accreting stars but is common to most young stars. Studying the abundances of pre-main sequence (PMS) stars and ZAMS stars together, we found that G-type stars show on average lower Ne/Fe and O/Fe abundance ratios than the M- and K-type stars. It seems therefore that abundances are a function of spectral type and are similar in PMS stars and in active main-sequence stars.

In conclusion, following the history of the coronal emission of the Sun in time we have found that the X-ray activity decreases from the youngest ages to the age of the present-day Sun. Younger stars rotate faster, and their internal dynamo is thus more efficient, resulting in increased magnetic activity. The X-ray luminosity and the coronal temperature are higher in ZAMS stars than in more evolved stars. The X-ray emission of non-accreting PMS stars is found to be consistent with the emission of ZAMS stars: All stars are saturated, and a relation between the coronal temperature and $L_X$ holds that is fully consistent with the equivalent relation in main-sequence stars. Further, PMS stars with $M \approx 1 M_\odot$ show $L_X$ and $T_{av}$ similar to ZAMS solar analogs ($L_X \approx 1 - 5 \times 10^{30}$ erg s$^{-1}$, log $T_{av}$[K] $\approx 7$). However, there is an appreciable change toward younger phases of a star’s life, when the star is still accreting material from the disk: Its X-ray emission is on average lower by a factor of $\approx 2$ and the
coronal temperatures are higher by a factor of $\approx 2.5$. At the same time, we find a soft excess in the high-resolution spectra of CTTS. We suggest a model in which accreting material cools the magnetic loops, so that the X-ray emission is shifted to cooler temperatures and becomes undetectable with CCD spectra. The X-ray emission at lower temperature results in soft emission that can be detected only in high-resolution spectra.

However, the number of pre-main sequence stars observed with high-resolution spectroscopy is still low. Further observations are needed to understand which properties are common to all CTTS.
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  A. Telleschi, M. Güdel, K. R. Briggs, M. Audard & L. Scelsi

- X-Ray Emission from T Tauri Stars and the Role of Accretion: Inferences from the XMM-Newton Extended Survey of the Taurus Molecular Cloud
  A. Telleschi, M. Güdel, K. R. Briggs, M. Audard, & F. Palla

- The First High-Resolution X-Ray Spectrum of a Herbig Star: AB Aurigae
  A. Telleschi, M. Güdel, K. R. Briggs, S. L. Skinner, M. Audard, & E. Franciosini

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- Some Like It Hot: The X-Ray Emission of the Giant Star YY Mensae
• A Deep Look at the T-Type Brown Dwarf Binary ε Indi Bab with Chandra and the Australia Telescope Compact Array
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• Evidence for an X-Ray Jet in DG Tauri A?

• The XMM-Newton Extended Survey of the Taurus Molecular Cloud
  (XEST)

• X-ray Emission from Jet-Driving Protostars and T Tauri Stars

• Unbinned Maximum-Likelihood Estimators for Low-Count Data. Applications to Faint X-Ray Spectra in the Taurus Molecular Cloud

• X-Ray Emission of the Young Brown Dwarfs of the Taurus Molecular Cloud

• The X-Ray Activity-Rotation Relation of T Tauri Stars in Taurus-Auriga

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- Statistics of Superimposed Flares in the Taurus Molecular Cloud
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- Spectral Properties of X-Ray Bright Variable Sources in the Taurus Molecular Cloud

**Conference Proceedings and Others**

- Coronal X-Ray Spectroscopy of Solar Analogs

- Coronal X-Ray Spectroscopy of Solar Analogs

- Coronal Evolution of Solar Analogs: A Study with XMM-Newton

- X-Ray Emission from Accreting, Jet-Driving T Tau Stars
  A. Telleschi, M. Güdel, K. Briggs, K. Arzner, S. Skinner, & M. Audard
  Protostars and Planets V, Posters, no. 8338

- XMM-Newton Survey of the Chamaeleon I Star Forming Region
  A. Telleschi, M. Güdel, K. Briggs, & M. Audard
List of Publications
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