X-RAY MODELING OF VERY YOUNG EARLY-TYPE STARS IN THE ORION TRAPEZIUM:
SIGNATURES OF MAGNETICALLY CONFINED PLASMAS AND EVOLUTIONARY IMPLICATIONS

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ABSTRACT

The Orion Trapezium is one of the youngest and closest star-forming regions within our Galaxy. With a dynamic age of $3 \times 10^7$ yr, it harbors a number of very young hot stars, which likely are on the zero-age main sequence (ZAMS). We analyzed high-resolution X-ray spectra in the wavelength range of 1.5–25 Å of three of its X-ray–brightest members ($\theta^1$ Ori A, C, and E) obtained with the High Energy Transmission Grating Spectrometer (HETGS) on board the Chandra X-Ray Observatory. We measured X-ray emission lines, calculated differential emission measure distributions (DEMs), and fitted broadband models to the spectra. The spectra from all three stars are very rich in emission lines, specifically from highly ionized Fe, which includes emission from Fe xvii to Fe xxv ions. A complete line list is included. This is a mere effect of high temperatures rather than an overabundance of Fe, which in fact turns out to be underabundant in all three Trapezium members. Similarly there is a significant underabundance in Ne and O as well, whereas Mg, Si, S, Ar, and Ca appear close to solar. The DEM derived from over 80 emission lines in the spectrum of $\theta^1$ Ori C indicates three peaks located at 7.9, 25, and 66 MK. The emission measure varies over the 15.4 day wind period of the star. For the two phases observed, the low-temperature emission remains stable, while the high-temperature emission shows significant differences. The line widths seem to show a similar bifurcation, where we resolve some of the soft X-ray lines with velocities up to 850 km s$^{-1}$ (all widths are stated as half-width at half-maximum), whereas the bulk of the lines remain unresolved with a confidence limit of 110 km s$^{-1}$. The broadband spectra of the other two stars can be fitted with several collisionally ionized plasma model components within a temperature range of 4.3–46.8 MK for $\theta^1$ Ori E and 4.8–42.7 MK for $\theta^1$ Ori A. The high-temperature emissivity contributes over 70% to the total X-ray flux. None of the lines are resolved for $\theta^1$ Ori A and E with a confidence limit of 160 km s$^{-1}$. The influence of the strong UV radiation field on the forbidden line in the He-like triplets allows us to set an upper limit on distance of the line-emitting region from the photosphere. The bulk of the X-ray emission cannot be produced by shock instabilities in a radiation-driven wind and are likely the result of magnetic confinement in all three stars. Although confinement models cannot explain all the results, the resemblance of the unresolved lines and of the DEM with recent observations of active coronae in II Peg and AR Lac during flares is quite obvious. Thus we speculate that the X-ray production mechanism in these stars is similar, with the difference that the Orion stars may be in a state of almost continuous flaring driven by the wind. We clearly rule out major effects due to X-rays from a possible companion. The fact that all three stars appear to be magnetic and are near zero age on the main sequence also raises the issue of whether the Orion stars are simply different or whether young massive stars enter the main sequence carrying significant magnetic fields. The ratio log $L_X/L_{bol}$ using the “wind” component of the spectrum is $-7$ for the Trapezium stars, consistent with the expectation from O stars. This suggests that massive ZAMS stars generate their X-ray luminosities like normal O stars and magnetic confinement provides an additional source of X-rays.

Subject headings: open clusters and associations: individual (Orion, Trapezium) — plasmas — stars: early-type — stars: formation — techniques: spectroscopic — X-rays: stars

1. INTRODUCTION

Since the discovery of X-ray emission from massive early-type stars more than two decades ago (Seward et al. 1979; Harnden et al. 1979), there has been an ongoing quest to explain its origins and to develop physical models that consistently predict its characteristics. Although to date no definitive models for the production of X-rays in stellar winds exist, it is widely thought that X-rays are produced by shocks forming from instabilities within a radiatively driven wind (Lucy & White 1980). This phenomenological model has been revised and expanded throughout the years. Lucy (1982) showed that these shocks can exist well out into the terminal flow, overcoming the attenuation problem of the previous model. Owocki, Castor, & Rybicki (1988) extended the model in that they showed that reverse shocks are much stronger than forward shocks in high-velocity gas at low densities, which they deduced from P Cygni absorption in UV resonance-line profiles (Puls, Owocki, & Fullerton 1993; Hillier et al. 1993). These models were able to successfully explain the mass loss from a hot luminous star in the UV domain as well as the soft X-ray spectral temperatures, but to date they cannot correctly predict observed X-ray fluxes. For example, Feldmeier, Puls, & Pauldrach (1997b) see the possibility of mutual collisions of dense gas shells in the outer wind producing stronger shocks. Claims that the X-ray emission could originate from coronal gas near the stellar photosphere were soon considered unlikely because of the lack of soft X-ray absorption edges (Cassinelli & Swank 1983) as well as the absence of coronal emission lines in optical spectra (Nordsieck, Cassinelli, & Anderson 1981).
That all O and early-type B stars are strong stellar X-ray emitters is now a well-established fact thanks to relentless observations with *Einstein* and *ROSAT* (Pallavicini et al. 1981; Chlebowski, Harnden, & Sciorinto 1989; Berghöfer & Schmidt 1994; Cassinelli et al. 1994). Typical X-ray luminosities are of the order of $10^{32}$ erg s$^{-1}$ for O stars and $10^{39.5}$–$10^{41.5}$ ergs s$^{-1}$ for B stars (Berghöfer, Schmitt, & Cassinelli 1996), while many low-mass (late-type) pre-main-sequence (PMS) stars radiate at orders of magnitude lower luminosity and only the peak of their luminosity function reaches $10^{31}$ ergs s$^{-1}$ (Feigelson & Montmerle 1999).

Because of the lack of spectral resolving power, however, many results from *Einstein* and *ROSAT* were based on statistical properties of a large sample of stars (Chlebowski et al. 1989; Berghöfer et al. 1996). Among these results were that the X-ray luminosity in early-type stars typically scales with the bolometric luminosity with log$(L_X/L_{bol}) = -7$ and, with a few exceptions, scales as high as $-5$.

In the first very detailed spectral analysis using the *ROSAT* PSPC, Hillier et al. (1993) fitted spectra of $\zeta$ Pup with non-LTE models under the assumptions that the X-rays arise from shocks distributed throughout the wind and that recombination occurs in the outer regions of the stellar wind. The best fits predicted two temperatures of $10^6$–$6.7 \times 10^6$ K with shock velocities around 500 km s$^{-1}$. On the basis of this approach, Feldmeier et al. (1997a) added the assumption that the X-rays originate from adiabatically expanding cooling zones behind shock fronts and described the spectra with postshock temperature and a volume filling factor. These results were also compared to results from Cohen et al. (1996), who used *ROSAT* and *EUVUE* data to constrain high-temperature emission models in the analysis of the B giant, $\epsilon$ CMa. A continuous temperature distribution was inferred over single- or even two-temperature models. The result of that comparison remained inconclusive, since both views appeared indistinguishable in the spectra. Despite the success of the wind-shock models, several unanswered issues remain from the *Einstein*, *ROSAT*, and *ASCA* era, which seem to be quite in contrast to this model (see also below). One issue concerns the unusually hard X-ray spectra of the B0.2 V star $\tau$ Sco observed with *ASCA* (Cohen, Cassinelli, & Waldron 1997), of $\lambda$ Ori (Corcoran et al. 1994), and of *ROSAT* spectra of stars later than B2 type (Cohen, Cassinelli, & MacFarlane 1997a).

With the availability of high-resolution spectra from *Chandra* and *XMM-Newton*, the spectral situation became much more complex and confusing. The first published high-resolution X-ray spectrum of the Orion Trapezium star $\Theta^1$ Ori C showed extreme temperatures and symmetric lines (Schulz et al. 2000, hereafter Paper I). These properties are not expected from shocked material near or beyond regions where the wind reached its terminal velocity. On the basis of a soft X-ray spectrum and symmetric emission lines from $\zeta$ Ori, Waldron & Cassinelli (2001) argued that the emitting plasma likely originates near the photosphere. Highly resolved spectra from $\zeta$ Pup with *XMM-Newton* (Kahn et al. 2001) and *Chandra* (Cassinelli et al. 2001) finally showed some expected emission characteristics, i.e., moderate temperatures of 5–10 MK and blueshifted and asymmetric lines. Such X-ray line profiles (Ignace 2001; Owocki & Cohen 2001) are significant characteristics of attenuated moving X-ray-emitting plasma. Schulz et al. (2001b) and Schulz (2002) report on similar evidence from line profiles in HD 206267 and $\iota$ Ori, respectively.

The fact that the Orion Trapezium star $\Theta^1$ Ori C shows such strange X-ray characteristics may not come as too much of a surprise, since this star was already known to be of a rather peculiar nature (Stahl et al. 1995; Gagne et al. 1997). Babel & Montmerle (1997a) proposed an aligned magnetic rotator model. The Orion Trapezium region was always quite difficult to observe prior to *Chandra* simply for the reason that its constituent members could never be spatially resolved. The *ROSAT* HRI (Gagne, Caillault, & Stauffer 1995) provided better spatial resolution but no spectral information. Yamauchi & Koyama (1993) observed hard X-rays from the Orion Nebula region with *Ginga* that did not rule out the possibility of $2$–$3$ keV X-rays from $\Theta^1$ Ori C but focused more on possible hard extended emission within the Nebula. This issue was studied further with *ASCA* (Yamauchi et al. 1996). Highly resolved images and spectra with *Chandra* could clearly resolve this issue. Schulz et al. (2001a) fully resolved the Orion Trapezium in the X-ray band between 0.1 and 10 keV and found no diffuse emission between the Trapezium stars. Furthermore, four of five of the brightest Trapezium stars showed hard X-ray spectra, with $\Theta^1$ Ori C being the hottest star, showing temperatures of up to $6 \times 10^6$ K (Paper I).

In this paper we further investigate these issues through detailed modeling of High Energy Transmission Grating Spectrometer (HEGTS) spectra from the three X-ray–brightest Trapezium stars $\Theta^1$ Ori A, C, and E (Schulz et al. 2001a). These stars are excellent candidates for such a common study since they are presumably very young. The median age of the Orion Trapezium stars is 0.3 Myr (Hillenbrand 1997), and they were born at the same time and should have had a quite similar initial chemical conditions. Their spectral types range from O6.5 V to B3.

2. CHANDRA OBSERVATIONS

We accumulate our spectra from two observations in the early phases of the *Chandra* mission. The first observation was performed on October 31 UT 05:47:21 1999 (OBSID 3) and lasted 50 ks. The second observation was obtained about 3 weeks later on 1999 November 24 UT 05:37:54 (observation ID [OBSID] 4) and lasted 33 ks. For more details of the observations and some of the analysis threads we refer to Paper I and Schulz et al. (2001a).

2.1. Data Analysis

We reprocessed the data using the most recently available calibration and CIAO implementations and produced event lists containing the proper grating dispersion coordinates. The spectral extraction was performed using CIAO tools, and we also used custom software for some of the broadband spectral analysis. The modeling of the spectra and its lines was done using ISIS, and the emission measure distribution was calculated as in Huenemoerder, Canizares, & Schulz (2001). As already described in Paper I, the Orion Trapezium is embedded in a cluster of fairly bright sources. We thus have to clean each spectrum from contributions of interfering cluster sources, which would imitate lines by

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1 See http://chandra.harvard.edu/CIAO2.3.

2 See http://space.mit.edu/ISIS.
coincidence, as well as from dispersed photons of other grating spectra crossing the dispersion track of interest. In our cases of interest the latter effect was entirely eliminated by the energy discrimination of the CCDs. Figure 1 shows the HETGS focal plane view near the zeroth order. In both exposures, at the top and bottom part of the figure we circled the zeroth-order positions of the main Trapezium stars. We also highlighted (for illustration purposes only) the tracks of the dispersed spectra for the brightest source θ1 Ori C.

Many spectral tracks from various Trapezium cluster stars are detected together with point sources at different off-axis positions that scatter throughout the entire array. One of the differences from the extraction procedure outlined in Paper I is that we now have to also extract the spectra of θ1 Ori A and E, which have a separation of a few arcseconds and are thus more likely to interfere with each other. We make extensive use of the fact that we have observations at two different roll angles, which gives very good separation at least at one roll angle. Besides visual inspection, we also use the fact that we know the relative fluxes of the two sources from the zeroth-order CCD spectra. In order to minimize the effect of interfering point sources and parallel dispersion tracks of close sources, we sometimes narrowed the cross-dispersion selection range. The standard selection range in cross-dispersion engulfs 95% of the flux, and in cases where we narrowed this range we have to renormalize the fluxes. However, we note that by correcting for the expected cross-dispersion profile, we add some systematic errors once we adjust final fluxes. We also correlated detected source point-spread functions with the dispersion tracks of the three targets and eliminated the data in the case of a true interference.

2.2. Raw and Exposure-corrected Spectra

Figure 2 shows the first-order HEG spectra of θ1 Ori C for both observation periods after all cleaning cycles. The weaker appearance of the 1999 November spectra compared to the 1999 October spectra is mostly due the difference in exposure, but for some lines there may also be variability related to the 15.4 day wind cycle. Here the October observation corresponds to phase 0.82 and the November observation to phase 0.37, using the ephemeris from Stahl et al. (1996). In order to investigate this effect we added spectra from the MEG and HEG and divided the exposure-corrected flux spectra of the two phases. This spectral ratio is shown in Figure 3. We smoothed and binned the spectra into large (0.5 Å) bins. The ratio shows that there is a flat part about 38% above unity below 6 Å. The detailed analysis below shows that the changes are due to variable emissivity at high temperatures. The mean (rms) difference across the whole band is about 24%. Schulz et al. (2001a) report different fluxes in the zeroth of about 10%, which is consistent given the uncertainties in that observation due to a pile-up fraction of over 20%.

2.3. Line Widths

A controversial issue in Paper I (see Schulz et al. 2003) was the analysis of the line widths in the case of θ1 Ori C. Because of a software error the line width presented in that paper did not properly account for the response of the instrument and, specifically, underestimated the amount of line blends between 10 and 12 Å. Thus the lines appeared wider than they really are. Here we present an analysis that properly includes the instrument response and takes care of all the line blends. The lines below about 13 Å clearly appear unresolved. Figure 4 (top) shows lines from three high-energy H-like ions, from S xvi, Si xiv, and Mg xii. We chose these lines because they are most likely not affected by blends. The model used for the line fits is described in § 3.2.2, and it fits not only the local but the overall broadband continuum. Thus the model for all three lines comes from the same fit (Fig. 4, bottom). Shown are the co-added spectra and models (red curves). In addition we show a broad stretch from the same model fit between 10.5 and 12 Å showing many highly ionized and blended Fe states. The model sufficiently describes these blends with no intrinsic broadening. For unresolved lines we can set a 90% confidence limit of the half-width at half-maximum (HWHM) of 110 km s⁻¹ on average based on the statistical properties of the spectrum.

There is a second population of lines at longer wavelengths, which are resolved. The statistical quality for many of them is marginal as in this bandpass the spectrum is effectively absorbed. Figure 5 shows the two brightest resolved lines, one from an Fe xvii ion at 15.01 Å and the O viii line...
at 18.97 Å. The Fe xvii can be fitted with a HWHM of 460 km s\(^{-1}\), and the O viii line can be fitted with a HWHM of 850 km s\(^{-1}\). The latter is also the only one we find slightly blueshifted by about 240 km s\(^{-1}\). In general it is expected to resolve lines better at higher wavelengths since the FWHM is constant with wavelength. However, these velocities exceed the confidence limit of 110 km s\(^{-1}\) for the unresolved lines considerably, and this indicates intrinsic broadening for some lower ionization states.

### 3. SPECTRAL ANALYSIS

We divide the analysis of the \(\Theta^1\) Ori C spectrum into two parts: we model the broadband spectrum with various temperature components of hot plasmas, and we construct an emission measure distribution from single-line emissivities. For the latter we need a large number of significant line fluxes, which we have available only in the case of \(\Theta^1\) Ori C. For the other hot stars we then compute variations of the

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**Fig. 2**—HEG first-order spectra of \(\Theta^1\) Ori C. The top two panels show the +1 and −1 spectrum of the first observation, and the bottom two panels show the same for the second observation.
broadband plasma model derived from the \( \Theta^{1} \) Ori C spectrum.

### 3.1. The Spectral Model

For the spectral modeling we exclusively use the Astrophysical Plasma Emission Code and Database (APEC and APED),\(^3\) described by Smith et al. (2001). The database is available in ISIS, and we can compute emissivities for a collisionally ionized equilibrium plasma in terms of temperature, density, and various abundance distributions. In Paper I we deduced that these spectra require a range of temperatures. For the model we assume that all emitting plasmas contributing to the spectrum are in collisional ionization equilibrium. This means that a thermal plasma is in a stable ionization state under the coronal approximation. In this approximation it is assumed that the dominant processes are collisional excitation or ionization from the ground state balanced by radiative decay and recombination. Photoionization and photoexcitation as well as collisional ionization from excited states are assumed to be negligible. Note that because of the usually quite strong UV radiation field near the stellar surface, this assumption is not quite true and the metastable forbidden lines in the He-like triplets may be affected. We will discuss the He-like triplets in a separate section.

From over 80 significant lines we calculate a differential emission measure (DEM) distribution for the two phases separately and for the phase-averaged spectra. In the other two stars, \( \Theta^{1} \) Ori A and E, we lack this large number of lines and cannot calculate such a significant DEM. We therefore approximate the phase-averaged DEM of \( \Theta^{1} \) Ori C by a few temperature components, for which we compute model spectra. We allow as many temperature components as necessary to account for 95% of the lines (3 \( \sigma \)) observed. Once adding temperature components does not improve the fit, we allow the abundance distribution to vary against a solar element abundance distribution. A similar procedure was followed to also fit the spectra obtained for \( \Theta^{1} \) Ori A and E.

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\(^3\) See http://hea-www.harvard.edu/APEC.
ion states larger than Fe\textsuperscript{xxi}, specifically Fe\textsuperscript{xxiv}, and ion states lower than Fe\textsuperscript{xix}, specifically Fe\textsuperscript{xvii}. The fractional abundances (i.e., the ratio of integrated flux of a specific ion species over all Fe emission) show a dominance of Fe\textsuperscript{xxiii} to Fe\textsuperscript{xxv} ions (0.36), as well as Fe\textsuperscript{xvii} to Fe\textsuperscript{xix} (0.58), but a significant lack of Fe\textsuperscript{xx} to Fe\textsuperscript{xxii} ions (0.06). The line positions agree well within the expected uncertainties with the positions produced by APED. The second part of Table 1 shows the list of ions with Z lower than 26. The results are very similar to what we observe for Fe ions, i.e., lines from low and high temperatures dominate the fractional abundances.

Figure 7 (top) shows the DEM versus temperature over a range 3–300 MK. The middle (red) line shows the distribution, the upper and lower lines (green) trace the 90% uncertainty range. There are clearly three peaks visible at 7.9 ± 0.2, 26.3 ± 1.8, and 66.1 ± 11.2 MK. The emissivity is dominated by plasmas at temperatures higher than 15 MK with a temperature tail that allows the distribution to exceed 100 MK. There is a
prominent gap at 7.15 MK that divides the DEM into emission from high- and low-temperature plasmas. This drop roughly corresponds to the missing emissivities from Fe\textsuperscript{xx} to Fe\textsuperscript{xxii} ions in Table 1. The low-temperature peak incorporates most ions from O to Si and Fe\textsuperscript{xvii} to Fe\textsuperscript{xxii}, while the high-temperature DEM corresponds to Fe\textsuperscript{xxiii} to Fe\textsuperscript{xxv}, Si to Ca. There is also an incision in the high-temperature part of the DEM that roughly corresponds to the weak showing of S\textsuperscript{XVI} and Ca\textsuperscript{xix} in one of the spectra (see below).

The middle part of Figure 7 shows the DEMs for two phases separately. The solid lines show the upper and lower error limits of the DEM for phase 0.82, and the dotted lines show the same for phase 0.37. There are a few remarkable characteristics. The incision between high and low temperature at 7.15 MK persists in both phases, and the low-temperature emissivity is similar. The high-temperature emissivity shows a significant difference in that the middle peak is much more pronounced in phase 0.82, whereas the high-temperature peaks appear unchanged. This means the overall variability in line emissivity is predominately in the Fe\textsuperscript{xxii} and Fe\textsuperscript{xxiv} lines as well as in the Si\textsuperscript{xiv}, where we observe the bulk of the emissivity.

The DEM fit also adjusts abundances (Fig. 7, bottom) with respect to solar abundances. Here we find for Mg, Si, S, Ar, and Ca ions no significant deviation from solar values. We find for O a factor of 0.2 ± 0.1 for Ne and for Fe a factor of 0.5 ± 0.1 underabundance. We find this distribution in both observations. Additional uncertainties in these values may stem from uncertainties in the ionization balance (Mazzotta et al. 1998). However, we also have to stress the point that once we apply the spectra for each ion fitted
The determination of the continuum in the DEM analysis is an iterative process assuming that all continuum contributions come from the same thermal plasma generating the emission lines. Contributions from a nonthermal component critically affect the Fe abundances. A predicted shape of $\propto E^{-1/2}$ of a Compton component (Chen & White 1991) most prominently contributes to the continuum near the Fe xxv line. However, the continuum level there is already quite determined by the fits to the Fe xxiv lines and their adjacent continuum levels. We estimate that the contribution cannot be more than about 1/2 to the total X-ray flux.

### 3.2.2. Phase-averaged Plasma Model

Figure 8 shows the phase-averaged count spectrum binned by 0.005 Å. MEG and HEG have been added. In this approach we model this spectrum with a few constrained components by approximating the emission measure above. This section in this regard does not produce much new information, but it verifies the method used for the fainter stars. However, for reasons of consistency we use the luminosities and fluxes determined in this section for further discussion. We rebin the above DEM into a few coarse intervals and calculate a model component for each DEM interval. We could calculate the model directly from each DEM bin, i.e., from the log $T/K = 0.05$ grid. However, the idea is to produce a simplified model in good approximation to the DEM. The models were again calculated using the APED database, folded through the spectral response function, and then fitted to the measured spectra.

We fitted only one component at a time (keeping the parameters of the other ones fixed) and repeated this step with the other components until the final spectrum meets the criteria below. To sensitively constrain the model, we use the DEMs above, which helps us to constrain the relative strengths of the components, i.e., the level of the continuum, the position of its high-energy cutoff, the strength of the Fe xxv lines, and various temperature-dependent line ratios.
range 1–25 Å agree within 2%, the continuum is well represented, and the line ratios agree within 10%.

The final result is overplotted in Figure 8. The thermal flux continuum shows a cutoff around 3.4 Å. In order to get this cutoff we need a high temperature of 66.5 MK. The shallow decline of the continuum below the cutoff requires stronger intermediate temperature components (43.8 and 29.5 MK), which are constrained by line ratios of Si, S, and Fe lines as well as the fact that the sum of the Fe xxv line fluxes of these two high-temperature components need to match the observed line flux. These three components account for almost all of the continuum, providing 75% of the star’s X-ray luminosity. The low-temperature components at 6.1, 9.3, and 16.5 MK produce many lines between 9 and 20 Å but account for only a fraction of the total luminosity. In fact, if the star did not have hot components, it would be faint. From Table 2 the X-ray luminosity accounting for the low-temperature DEM component only is $3.5 \times 10^{31}$ ergs s$^{-1}$. Using the bolometric luminosity as listed by Berghofer et al. (1996), we find $-7.2$ for the low-temperature log $L_X/L_{bol}$. More than 85% of the luminosity is radiated by plasmas with temperatures higher than $1.5 \times 10^7$ K, and 75% is from temperatures higher than $3 \times 10^7$ K. This is consistent with the results from the zeroth-order CCD spectra (Schulz et al. 2001a). The abundances had to be adjusted during the fit as well, and the result is similar to the distribution observed in the DEM analysis.

3.2.3. Optical Depths, Formation Radii, and Densities

Line ratios may be used to place limits on optical depths, line formation radii, and densities. Schmelz, Saba, & Strong (1992) suggested that the Fe xvi (upper-level 2p$^5$3d$^1P_1$) line at 15.01 Å could be used as a measure for resonant scattering (and thus density) in the emitting plasma. Specifically, the ratio to its Fe xvi resonance-line neighbor at 15.26 Å ($2p^6 \, ^1S_0-2p^53d^1D_1$) is of interest (see also Waljeski et al. 1994, Brickhouse et al. 2000, and a discussion in Huenemoerder et al. 2001). From the APED database we deduce a theoretical ratio of 3.57 for the optically thin limit. The measured ratio in the spectrum is $3.47 \pm 0.12$, and thus despite the large error bar indicates agreement with a low-density plasma. The formation radius of this line can be estimated from its line width. This line was indeed resolved with a HWHM of 460 km s$^{-1}$. The standard law of velocity for the acceleration zone of the wind $v(r) = v_\infty (1 - r_{\text{star}}/r)^{\beta}$ (Lamers & Cassinelli 1999) with the index $\beta = 0.88$ and a terminal velocity of 1000 km s$^{-1}$ would then locate the line-emitting region to slightly more than half a stellar radius from the photosphere.

Another line that was resolved was the O viii line at 18.97 Å. Its HWHM corresponds to velocity of 850 km s$^{-1}$, which would place the emitting region almost near the terminal velocity of the wind at about 7 stellar radii from the photosphere. In contrast, the line width limit of 110 km s$^{-1}$ would place the emitting radius of all the other (non–He-like) lines to within 10% of a stellar radius above the photosphere (see also Cohen et al. 2002 for τ Sco). Whether there should be an asymmetry of the line due to occultation seems irrelevant since we do not resolve most of the lines.

The forbidden lines in the He-like triplets are metastable, and their ratio with the corresponding intercombination lines is density sensitive above some ion-dependent threshold. Figure 9 shows the observed He-like triplets from S xv,
Fig. 8.—Measured phase-averaged count spectrum over the full exposure for θ² Ori C. The bins of the spectrum have a value of 0.005 Å. The red line is the approximated phase-averaged plasma model deduced for this spectrum.
Si xii, and Mg x. The model fits for these triplets have been optimized for each element specifically to fit the resonance line. One of the most striking effects seen in the triplets from O vii (very faint), Ne ix, and Mg x (Fig. 9, bottom) is that the forbidden line is not observed, while it still appears quite prominent in Si xii and S xv (Fig. 9, middle and top). Although the forbidden transition in the Si xii triplet is blended with the Lyγ line of Mg xii, based on the flux observed in the Mg xii Lyα line, the Lyγ contribution cannot be more than 5%. However, we observe a forbidden-line flux comparable to the expectation of the model. We observe a similar picture in S xv (which is not blended); however, the $f_i$/$i$ ratio here is not sensitive to densities lower than $10^{14}$ cm$^{-3}$.

All the triplets are subject to UV excitation, and it has been shown by Kahn et al. (2001; see also Blumenthal, Drake, & Tucker 1972) in the example of ζ Pup that the low $f_i$/$i$ ratios in line-driven wind plasmas are not likely an effect of collisional excitation in high-density plasmas but rather are due to the destruction of the forbidden line by the large UV flux at the corresponding excitation wavelengths. Waldron & Cassinelli (2001), Cassinelli et al. (2001), and Miller et al. (2002) use model UV fluxes (e.g., Chavez, Stalio, & Holberg 1995) to estimate radial constraints on the X-ray-emitting plasma. Such an analysis in the case of θ1 Ori C is highly delicate. One critical item is the modeling of the actual UV flux between ~900 and 1500 Å, which provides the source for the photoexcitation in Si xii and Mg x. We checked the ratio of the UV spectra of θ1 Ori C and ζ Pup using Copernicus data (Snow & Jenkins 1977), and after correcting for the different stellar parameters (19 R$_\odot$, 42,500 K, $E_{B-V} = 0.07$ for ζ Pup; 8 R$_\odot$, 39,000 K, $E_{B-V} = 0.35$ for θ1 Ori C) (Pauldrach et al. 1994; Howard & Prinja 1989; Berghöfer et al. 1996), we could not quite reconcile the model with the data in a sense that the UV flux in θ1 Ori C seems lower than the model. A proper correction for extinction is certainly critical.

Effectively there is no obvious reason why the UV field should be weaker than one would expect from blackbody model atmospheres (MacFarlane et al. 1993; Chavez, Stalio, & Holmberg 1995). In this respect we estimated the formation radius by assuming a blackbody spectrum of 39,000 K for the radiation field for an O7 V star (Howard & Prinja 1989) as an upper limit. The surface temperature of θ1 Ori C is also a source of uncertainty, as θ1 Ori C is classified somewhere between O6 and O7.5, sometimes even as O4. In the case of the latter the result can differ by over 60%. We calculated the photoionization (PE) rates using the recipe provided by Kahn et al. (2001; see also Mewe & Schrijver 1978) and obtained 2 $3S-1S$ decay rates from Drake (1971). In Mg x the forbidden line can be only marginally detected above the 1 σ error of the continuum, and here we consider the radius where the PE rate is of the order of the decay rate as critical. The $f_i$/$i$ ratios in Si and S are, although reduced from the one expected from the atomic data in the optically thin limit, significantly larger. Here we have to assume that the formation radius is farther away from the stars surface and that the PE rate is correspondingly smaller. Thus we find an upper limit to the formation radius of 4.2 stellar radii for Mg, 2.0 stellar radii for Si, and 1.3 stellar radii for S.

On the other hand it should be mentioned that in the case of zero UV flux the $f_i$/$i$ ratio of the triplets can also be used as density diagnostics. We then obtain densities of less than $4 \pm 2 \times 10^{13}$ cm$^{-3}$ for Mg and less than $9 \pm 8 \times 10^{13}$ cm$^{-3}$ for Si, repeating the result stated in Paper I.

### 3.3. θ1 Ori E and A

The next two X-ray brightest stars in the Trapezium cluster are stars A and E, with X-ray fluxes of $(1.3-2.5) \times 10^{-12}$ ergs cm$^{-2}$ s$^{-1}$ corresponding to luminosities of $(2.4) \times 10^{31}$ ergs s$^{-1}$ (Schulz et al. 2001a). These fluxes are an order of magnitude fainter than θ1 Ori C, which is likely due to their B spectral types (Cassinelli et al. 1994). In this respect the HETGS spectra are less brilliant; however, in both spectra we are still able to detect quite a number of emission lines and strong continua. This at least allows us to sufficiently constrain our multicomponent model. Table 3 shows the characteristics of the brightest lines for both stars. These line properties are very similar to the characteristics exhibited by θ1 Ori C. This is specifically noteworthy for the

### Table 2

| Star          | Component | $\log T$ (K) | Norm (ergs s$^{-1}$ cm$^{-2}$) | $L_X$ (1–10 keV) (ergs s$^{-1}$) |
|---------------|-----------|--------------|-------------------------------|----------------------------------|
| θ1 Ori C      | 1         | 6.79 ± 0.06  | 6.55 $\times 10^{-13}$        | 7.93 $\times 10^{20}$           |
|               | 2         | 6.97 ± 0.07  | 2.22 $\times 10^{-12}$        | 2.69 $\times 10^{21}$           |
|               | 3         | 7.22 ± 0.07  | 1.24 $\times 10^{-12}$        | 1.50 $\times 10^{21}$           |
|               | 4         | 7.47 ± 0.08  | 4.62 $\times 10^{-12}$        | 5.60 $\times 10^{21}$           |
|               | 5         | 7.64 ± 0.06  | 2.51 $\times 10^{-12}$        | 3.04 $\times 10^{21}$           |
|               | 6         | 7.82 ± 0.07  | 5.20 $\times 10^{-12}$        | 6.30 $\times 10^{21}$           |
| θ1 Ori E      | 1         | 6.63 ± 0.08  | 1.55 $\times 10^{-13}$        | 1.87 $\times 10^{20}$           |
|               | 2         | 7.02 ± 0.08  | 3.78 $\times 10^{-13}$        | 4.60 $\times 10^{20}$           |
|               | 3         | 7.31 ± 0.07  | 4.88 $\times 10^{-13}$        | 5.92 $\times 10^{20}$           |
|               | 4         | 7.49 ± 0.06  | 1.85 $\times 10^{-13}$        | 2.23 $\times 10^{20}$           |
|               | 5         | 7.67 ± 0.06  | 1.31 $\times 10^{-12}$        | 1.59 $\times 10^{21}$           |
| θ1 Ori A      | 1         | 6.68 ± 0.09  | 3.41 $\times 10^{-14}$        | 4.13 $\times 10^{29}$           |
|               | 2         | 7.06 ± 0.08  | 3.54 $\times 10^{-13}$        | 4.29 $\times 10^{29}$           |
|               | 3         | 7.34 ± 0.08  | 1.08 $\times 10^{-13}$        | 1.31 $\times 10^{30}$           |
|               | 4         | 7.48 ± 0.07  | 4.28 $\times 10^{-13}$        | 5.18 $\times 10^{30}$           |
|               | 5         | 7.63 ± 0.07  | 2.76 $\times 10^{-13}$        | 3.35 $\times 10^{30}$           |
components 2–5 in higher temperature seem similar to com-
ponents 2–5 in higher temperature. In this respect we observe a similar

temperature missing. In this respect we observe a similar

velocities near the onset of the wind cannot produce enough

most He-like triplet are too weak to perform a meaning-
ful analysis except for Si xiii, where in the case of star E
(Fig. 10, 6.74 Å) we observe a forbidden line that is almost
as strong as the resonance line. In star A it is not as strong
but clearly detectable. In both cases we do not detect an
intercombination line, and by setting a limit to the line using
the 1 σ error of the continuum we can get only a rough esti-
mate of the f/i ratio. We nevertheless can put some limits
on the formation radius of Si xiii. In the optically thin limit
without photoexcitation we expect 2.16 for the f/i ratio.
For star E we can limit the ratio to 3.43 ± 1.69, consistent
with the optically thin limit, and basically no UV destruc-
tion. Assuming a B0.5 V star with a surface temperature of
23,000 K, we would expect the forbidden line would get
departed completely within 0.4 stellar radius above the
photosphere. Since this is not the case and the line seems to
be fully intact, we assume that the 2 3S–1S decay rate
entirely dominates the PE rate, which should occur at a dis-
tance of 1.4 stellar radii. For star A the ratio is 2.13 ± 1.33,
and we can make the same argument, assuming a surface
temperature of 20,000 K (B1 V; Loyd & Strickland 1999). In
terms of density, i.e., under the assumption that there is no
photoexcitation, the f/i ratios seem to be in compliance
with the densities below 10^{13} g cm^{-3}.

4. DISCUSSION

High-resolution X-ray spectra provide new and more
powerful diagnostics of the high-energy emission from hot
star plasmas. These diagnostics involve line identifications,
the relation of specific line ratios with physical parameters
such as temperature, density, line shapes and shifts, as well
as full-scale plasma models. In Paper I we presented a pre-
liminary line list for θ¹ Ori C. For this paper we added one
more observation to the analysis for a net exposure of 83 ks.
The availability of improved calibration products and mod-
els allowed us to refine these results in major areas. For
θ¹ Ori C the number of detected lines increased and allowed
for better constraints on the plasma modeling specifically
for the calculation of the DEM. The DEM analysis of the
phase-averaged spectrum as well as the fitting of the broad-
band plasma model showed that the bulk of the X-ray emis-
sion comes from plasmas with temperatures above 15 MK,
and only a small, well-separated part of the emission
corresponds to emission of lower than 10 MK.

There are models related to the standard line-driven–
wind instability model that are able to allow temperatures in
excess of 10 MK (Feldmeier et al. 1997a; Howk et al. 2000;
Runacres & Owocki 2002). However, possible high shock
velocities near the onset of the wind cannot produce enough
volume emissivity to account for the X-ray spectrum. The results also confirm a similar behavior for θ¹ Ori A and E. This means the bulk of the X-ray emission in the bright Orion Trapezium stars is not compatible with any form of instabilities in a line-driven wind. In fact, these stars seem to be of a hybrid nature where only a small fraction of the X-rays are produced in wind shocks; a larger fraction shows a magnetic origin and bears striking similarities to the hard X-ray emission pattern observed in stars with active coronae (Huenemoerder et al. 2001, 2003).

The DEM model allows for a nonthermal continuum component that contributes less than 1% to the total flux for a power-law index of −0.5, as suggested by Chen & White (1991). These authors proposed a nonthermal origin for possible hard X-rays in form of an inverse Compton continuum. This limit corresponds to $4 \times 10^{30}$ ergs s⁻¹, which is an order of magnitude higher than predicted. Thus we may not really be sensitive to the issue. It is quite clear though that the high-energy flux from the Orion Trapezium stars is not due to a possible nonthermal origin.

### 4.1. Possible Shocked-Wind Component

The emissivity peak at log $T = 6.9$ includes most lines from O to Si and ions below Fe xx. These are the moderate temperature lines that we also observe in wind shocks in ζ Pup and others (Cassinelli et al. 2001; Kahn et al. 2001; Waldron & Cassinelli 2001; Schulz et al. 2001b; Schulz 2002), suggesting a similar origin for this component on θ¹ Ori C. This peak does not change within the two observed phases of the 15.4 day cycle in θ¹ Ori C. Ignace (2001) and Owocki & Cohen (2001) calculated the shape of X-ray line profiles in stellar winds under various conditions. The fact that most lines in the spectra do not show significant shifts and are unresolved puts limits on these conditions. If the lines were to be produced at large radii in the outer wind, we should observe broad, symmetric, sometimes flat-topped lines. Yet for wind shocks, the only way to achieve symmetry is to have almost no attenuation in wind. Given the low mass-loss rate of $4 \times 10^{-7} M_\odot$ yr⁻¹ (Howarth & Prinja 1989), this is quite likely. Some of the X-ray–emitting plasma also has to be near the photosphere at the onset of the wind, given the unresolved nature of the lines. However, it seems that some of the softest lines are indeed resolved and show velocity broadening of up to 850 km s⁻¹. This is indicative that these lines contributing to the low-temperature emissivity peak are due to X-rays from shocks in the outer wind. Would this low-temperature emissivity be the only contribution to the X-ray emission, the Orion stars would be quite X-ray faint, but with a log $L_X/L_{bol}$ between −7.2 and −7.6, which is near the canonical value for O stars.

### 4.2. Magnetically Confined Winds

The truth is, however, that log $L_X/L_{bol}$ is more of the order of −6.5 for these stars. Since most emissivity appears to originate at temperatures that are incompatible with

### Table 3

Identified Ions in θ¹ Ori A and E

| Ion      | $\lambda_0$ (Å) | Flux (θ¹ Ori A) (10⁻⁶ photons s⁻¹ cm⁻²) | Flux (θ¹ Ori E) (10⁻⁶ photons s⁻¹ cm⁻²) |
|----------|-----------------|---------------------------------------|---------------------------------------|
| Fe xxv   | 1.861           | ...                                   | 7.182 ± 1.077                         |
| Ca xix   | 3.198           | 0.895 ± 0.134                         | 1.481 ± 0.222                         |
| Ar xvii  | 3.734           | ...                                   | 3.321 ± 0.498                         |
| Ar xvii  | 3.949           | 0.298 ± 0.045                         | 2.274 ± 0.341                         |
| Si xvi   | 4.730           | 0.840 ± 0.126                         | 5.233 ± 0.785                         |
| Si xvii  | 5.039           | 1.097 ± 0.165                         | 2.737 ± 0.411                         |
| Si xviii | 5.217           | ...                                   | 1.194 ± 0.179                         |
| Si xiv   | 6.183           | 3.293 ± 0.494                         | 3.935 ± 0.590                         |
| Si xiii  | 6.648           | 1.731 ± 0.260                         | 2.573 ± 0.386                         |
| Si xii   | 6.740           | 1.706 ± 0.256                         | 2.751 ± 0.413                         |
| Fe xxiv  | 7.989           | 0.676 ± 0.101                         | ...                                   |
| Mg xii   | 8.420           | 2.612 ± 0.392                         | 5.060 ± 0.759                         |
| Mg xi    | 8.815           | 1.363 ± 0.204                         | ...                                   |
| Ne x     | 9.169           | 1.963 ± 0.294                         | 2.495 ± 0.374                         |
| Ne x     | 9.708           | 0.811 ± 0.122                         | 2.393 ± 0.359                         |
| Ne x     | 10.239          | 3.078 ± 0.462                         | 3.530 ± 0.530                         |
| Fe xxiv  | 10.619          | 2.919 ± 0.438                         | 3.508 ± 0.526                         |
| Fe xxii  | 10.983          | ...                                   | 2.048 ± 0.307                         |
| Fe xxv   | 11.176          | 1.013 ± 0.152                         | 4.392 ± 0.659                         |
| Fe xxv   | 11.376          | 2.186 ± 0.328                         | 2.875 ± 0.431                         |
| Fe xviii | 11.330          | ...                                   | 3.442 ± 0.516                         |
| Ne x     | 12.135          | 14.589 ± 2.188                        | 16.331 ± 2.450                        |
| Fe xvii  | 12.261          | 1.558 ± 0.234                         | 0.498 ± 0.075                         |
| Fe xx    | 12.576          | ...                                   | 1.617 ± 0.242                         |
| Ne x     | 13.445          | 5.727 ± 0.859                         | 37.722 ± 5.658                        |
| Fe xvii  | 15.014          | 21.573 ± 3.236                        | 4.164 ± 0.625                         |
| Fe x    | 15.079         | 8.305 ± 1.246                         | ...                                   |
| O viii   | 15.176          | ...                                   | 18.360 ± 2.754                        |
| O xvii   | 17.051          | 5.309 ± 0.796                         | 13.059 ± 1.959                        |
| O x     | 18.970          | 0.177 ± 0.027                         | 9.362 ± 1.404                         |
Fig. 10.—Measured count spectrum for θ^1 Ori E. The red line is the best fit found using a multitemperature APED model.
Fig. 11.—Measured count spectrum for 2 Ori A. The red line is the best fit found using a multitemperature APED model.
wind shocks, we have to look for other mechanisms. There are several indications that the enhanced X-ray activity could be triggered by magnetic fields. Gagne et al. (1997) interpreted the strong 15.4 day period in the X-ray and optical emission of θ1 Ori C reported by Stahl et al. (1996) in terms of the star’s significant magnetic field.

Babel & Montmerle (1997a) proposed a magnetically confined wind shock (MCWS) model based on an oblique magnetic rotator model for θ1 Ori C. Here the wind is confined by a magnetic dipole field and forced into the magnetic equatorial plane. The observed 15.4 day variability is thus produced by the tilt of the dipole field relative to the rotational axis. The predicted field is of the order of 300 G. Very recent spectrophotometric observations indicate a dipole field of 1.1 kG with an inclination of 42° with respect to the stars rotational axis (Donati et al. 2002). These values are quite consistent with the temperatures that we observe in the spectrum of θ1 Ori C. This model successfully explains the periodicity in the Hα and Hβ lines as well as in P Cygni line profiles in the UV (Stahl et al. 1993, 1996; Walborn & Nichols 1994; Reiners et al. 2000). In the X-ray light curve this interpretation of the periodicity is also attractive.

The HETG spectra, however, indicate a more complex behavior. The HETG observations were performed at phases around 0.37 and 0.82. According to the ROSAT HRI light curve in Gagne et al. (1997), we should see a difference in flux between the two observations of about 30%. We applied the nonvarying low-temperature emissivity peak, and we see the rise toward the first high-temperature phase, which includes all the lines that were accessible to the ROSAT bandpass. The HETG data show that there is more going on than just a flux change. We observe a dramatic change of emissivity at log T(keV) = 7.4 between the two phases. Donati et al. (2002) point out that meaningful comparisons with predictions of the MCWS model can be achieved only from lines at extreme configurations (phases 0.0 and 0.5). Our phase should be somewhere in between these two extremes. The model states that the change in luminosity between the two phases above 2 keV (below ~6 Å) should decrease compared to below 2 keV (above ~6 Å). This is not what we observe. The difference below 6 Å is largest instead, with almost 40%. Furthermore, we see no obvious reason in the MCWS why the plasma temperature distribution should change in the way that we observe in the DEM.

More quantitative modeling including a more detailed energy balance treatment by Ud’Doula & Owocki (2002) allow us to further differentiate the phenomenology between a magnetic field and a stellar wind. By relating the magnetic energy density to the matter outflow in the wind a confinement parameter \( \eta \sim BR^2/M v_{\infty} \) can be defined relating the magnetic field density at a radius \( R \) from the star and the mass-loss rate \( M \) in a stellar wind of terminal velocity \( v_{\infty} \). Their MHD simulations showed that in the case of a large \( \eta \) value (~10) and closed magnetic field topologies near the star the wind is forced into looplike structures in which strong shock collisions produce hard X-rays. Ud’Doula & Owocki (2002) showed that these structures could generate enough emissivity at high temperatures by applying the standard shock jump condition from Babel & Montmerle (1997b). For observed values of magnetic field strength, terminal wind velocity, and mass-loss rates for θ1 Ori C, the \( \eta \) parameter is well over 10 (Gagne et al. 2001). Donati et al. (2002) applied the standard model devised by Babel & Montmerle (1997b) and found similar trends in terms of average temperature and density.

The looplike structures proposed by Ud’Doula & Owocki (2002) and Gagne et al. (2001) may also explain the structure we observe in the DEM distribution, where different peaks may relate to different confined structures. We can further speculate that changes at different phases refer to different structures. These confined structures may be quite unstable on short timescales, resulting in highly variable X-ray emission. Feigelson et al. (2002) recently detect rapid variability in the O7pe star G2 Ori A and suggested unseen companions for these stars (see also below). This may not be necessary. As an interesting analogy, the DEMs recently deduced from X-ray emission form active coronae in cool stars such as AR Lac (Huenemoerder et al. 2003) and II Peg (Huenemoerder et al. 2001) appear very similar during flares, where in addition to a small low-temperature peak one or more strong high-temperature peaks evolve. The nature of the spectra in these cases are of striking similarity to the ones that we observe in the Orion Trapezium with strong continua and unresolved lines. In this respect we may speculate that in the case of the Orion stars matter is constantly supplied into magnetic loops by the wind generating shock jumps on a permanent basis. In other words, these stars would be in a permanent state of flaring. The means of energy deposit is expected to be different in cool stars, which posses a dynamo and deposit energy into the plasma via reconnection. The mechanism in the winds of hot stars is more indirect via confining. Clearly, these details still have to be worked out.

If these shocks are magnetically confined, they are not likely to produce line shifts and significant line broadening. Donati et al. (2002) simulated dynamic X-ray line spectra for several model cases for θ1 Ori C based on the MCWS model and found narrow-line shapes at the observed phases. Any predicted line shifts are of the order of 150–200 km s\(^{-1}\) and below our resolving power. This is consistent with the line characteristics in all three Orion spectra.

### 4.3. Low-Mass Companions

The Trapezium stars are known to have one or more companions. Quite recently Weigelt et al. (1999) reported on the existence of a close, probably low-mass companion to θ1 Ori C and confirmed the companion detected by Petri et al. (1998) in θ1 Ori A. The companion in θ1 Ori C is as close as 33 mas. The close companion in θ1 Ori A has a separation of 202 mas. The only star in the Trapezium cluster with no detected companion is θ1 Ori E. On the basis of the median age of the cluster of 0.3 Myr (Hillenbrand 1997) and the ubiquity of nearby proplyds (O’Dell, Wen, & Hu 1993; Bally et al. 1998), which likely contain Class II T Tauri stars (Felli et al. 1993; McCaughrean & Stauffer 1994; Schulz et al. 2001a), we consider these companions to be young T Tauri stars. Weigelt et al. (1999) similarly suggest that the companion in θ1 Ori C is a very young intermediate- or low-mass (\( M < M_\odot \)) star on the basis of PMS evolutionary tracks.

The X-ray emission of most young PMS stars in the Orion Trapezium cluster is usually absorbed (Garmire et al. 2000), but in addition they also emit hard X-ray emission with temperatures of around 30 million K (Schulz et al. 2001a). Giant X-ray flares can reach up to 60–100 MK.
magnetic signatures, with \( \Theta^1 \) Ori C, A, and E by the fact that we do not observe variability in the light curve (here we do not take into account high-frequency variability observed by Feigelson et al. 2002). If we see contributions from the low-mass companion, it has to be persistent. X-ray luminosity functions from many star-forming regions peak at luminosities below 10^{31} \text{ergs s}^{-1}. This has also been observed for the low-mass PMS population in the vicinity of the Orion Trapezium (Schulz et al. 2001a; Feigelson et al. 2002). In this respect a major contribution from such a companion to the spectrum in \( \Theta^1 \) Ori C can be ruled out.

It is possible that X-rays from a low-mass companion make a significant contribution to \( \Theta^1 \) Ori A, since its flux components are an order of magnitude fainter. Even if there were such a contribution, it can be only be a small fraction of the total observed X-ray luminosity. This would be even more the case for \( \Theta^1 \) Ori E should it harbor an unseen companion. We can thus rule out that the X-ray emissivity pattern is due to a possible binary nature of the stars.

### 4.4. Evolutionary Implications

The very similar properties and morphologies in the spectra of \( \Theta^1 \) Ori A, C, and E raises an intriguing issue. With an ionization age of the nebula of about 0.2 Myr and a median age of the cluster of about 0.3 Myr, it is quite suggestive that the Trapezium stars are true ZAMS stars. Zero-age here is considered to be the time when energy generation by nuclear reactions first fully compensates the energy loss due to radiation from the stellar photosphere. In the case of \( \Theta^1 \) Ori C, it has been stated many times that the magnetic activity could possibly be of pre-main-sequence origin (Gagne et al. 2001; Donati et al. 2002). Gagne et al. (2002) noted that stars like \( \Theta^1 \) Ori C, or \( \tau \) Sco all show signs of magnetic activity and that they are all associated with young star-forming regions. With the Orion Trapezium we may have an indication that magnetic fields may be anticorrelated with age. Four of the five main Trapezium stars now show strong magnetic signatures, with \( \Theta^1 \) Ori A, C, and E being the most striking cases. So far the view of primordial high fields has been limited to peculiar (Ap and Bp) stars, and its has long been suspected that \( \Theta^1 \) Ori C is a candidate for an Op star (Gagne et al. 2001). \( \Theta^1 \) Ori D, so far also classified as a peculiar dwarf (B0.5 Vp), is weak in X-rays. The X-ray spectrum (Schulz et al. 2001a) does not indicate very high temperatures, and the grating data are too marginal to search for narrow lines. However, it seems that for some reason all Trapezium stars are chemically peculiar either by coincidence or because this peculiarity has something to do with magnetism and/or their young age.

In Table 3 we list some properties from presumably young massive stars. These properties include the spectral type, the assumed cluster age, star-forming region, and the currently known X-ray temperatures. The compilation is certainly far from complete, but it demonstrates that stars from regions younger than a million years show clear evidence for magnetic activity. The main identifier for magnetic activity is the extremely high (>10^7 K) persistent temperature. For the Orion stars and \( \tau \) Sco we additionally observe symmetric and narrow lines. HD 164492 is an interesting case as it is at the center of a very active star-forming region associated with the Trifid nebula, which very recently has been classified to be in a “pre-Orion” evolutionary stage (Leifoch & Cernicharo 2000) and thus may be the “youngest” massive ZAMS star in the sample. Although not fully resolved, its X-ray characteristics (Rho et al. 2001) as observed with ASCA seem similar to what has been observed for \( \Theta^1 \) Ori C with ASCA. Stars like HD 206267 and \( \tau \) CMa are at the center of more evolved clusters and do not show these characteristics. In fact, they behave very much like \( \zeta \) Pup (Kahn et al. 2001; Cassinelli et al. 2001) in that they show temperatures and line shapes consistent with shock instabilities in a radiation-driven wind (Wojdowski et al. 2002; Schulz et al. 2001b; Schulz 2002). Finally, \( \iota \) Ori and \( \zeta \) Ori are part of the Orion region that is older than the Trapezium cluster core and further evolved. Here we also see no indication for magnetic activity (Schulz 2002). \( \zeta \) Ori (Waldron & Cassinelli 2001) may be a special case in that it lacks high temperatures but possesses symmetric X-ray lines, which likely are due to a low opacity of the wind.

Table 4 may reflect an evolutionary sequence in which massive stars that enter the main sequence carry strong magnetic fields interacting with an emerging wind. In the table we list two types of X-ray luminosities based on the dichotomy observed in the Orion stars, \( L_X^{\text{mag}} \) for the amount produced by presumably magnetic confinement and \( L_X^{\text{wind}} \) for the amount from wind shocks. The bolometric luminosities for the \( L_X^{\text{wind}}/L_{\text{bol}} \) ratios were taken from Berghöfer et al. (1996). Once the star evolves, magnetic fields may become less important as mass-loss rates are higher. The X-ray emission sooner or later is fully generated by instabilities in the wind only. Here studies that search for magnetic fields in hot stellar winds (see Chesneau & Moffat 2002) are useful. An intriguing aspect on the Orion stars is that there also seems to exist relatively faint X-ray emission consistent with nonmagnetic wind shocks. Thus even if it is the case that most Orion stars are indeed just abnormal and strong magnetic fields are more the exception than the rule, then very young massive stars if not magnetic should at least scale with their bolometric luminosity with \( L_X^{\text{mag}}/L_{\text{bol}} \sim -7 \) and not higher. This may explain the relative X-ray weakness of HD 164492 A relative to \( \Theta^1 \) Ori C (J. Rho, private communication), although \( \Theta^1 \) Ori C may be even younger than HD 164492A. This may not explain the extremely weak X-rays flux from \( \Theta^1 \) Ori D though. From a projected bolometric luminosity for a B0.5 V star of 10^{38} \text{ergs s}^{-1} and an X-ray luminosity of 3 \times 10^{27} \text{ergs s}^{-1} (from Schulz et al. 2001a), we compute a \( L_X^{\text{wind}}/L_{\text{bol}} \) of ~8.5 for the star, which is very low. On the other hand we can turn the argument around and in the case of \( \tau \) Sco use \( L_X^{\text{mag}}/L_{\text{bol}} = -7 \) to estimate the amount of \( L_X^{\text{mag}} \) out of the observed ratio, as has been done in Table 4.

Too little is known about massive stars as they quickly evolve toward the main sequence. Although it is well accepted that many high-energy processes in young stellar objects (YSOs) are dominated by magnetic activity (see Feigelson & Montmerle 1999 for a review) and that hard X-rays in flares are produced by powerful magnetic reconstructions, star-disk activities, and jet formation, these findings are still confined to low-mass stars. In several models for Class I and II YSOs magnetic field play a crucial role in the X-ray production, in the form either of X-ray winds and

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4 See http://www.mssl.ucl.ac.uk/~gbr/rgs_workshop/workshop ADS_index.html.
accretion (Shu et al. 1997) or of accretion shocks (Hartmann 1998). It has yet to be established how these models for low-mass YSOs may relate to more massive stars.

One should be careful, however, to simply argue with age in such a correlation. It has to be kept in mind that there are similar patterns once we include mass-loss rate and wind velocities in such a correlation. It has to be kept in mind that there are low-mass YSOs may relate to more massive stars.

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### REFERENCES

Babel, J., & Montmerle, T. 1997a, ApJ, 485, L29

Bally, J., Sutherland, R. S., Devine, D., & Johnstone, D. 1998, AJ, 116, 293

Berghöfer, T. W., & Schmitt, J. H. M. M. 1994, A&A, 292, L5

Bally et al. 1998

Bally et al. 1997b

Babel, J., & Montmerle, T. 1997a, ApJ, 485, L29

Cassinelli, J. P., & Swank, J. H. 1983, ApJ, 271, 681

Casinelli, J. P., Cohen, D. H., Cassinelli, J. P., Cohen, D. H., Cassinelli, J. P., & Waldron, W. L. 1997b, ApJ, 488, 397

Cassinelli, J. P., MacFarlane, J. J., Sanders, W. T., & Welsh, J. J. 1997a, ApJ, 487, 867

Cassinelli, J. P., Owocki, S. P., & Liedahl, D. A. 2003, ApJ, 586, 495

Chen, W., & White, R. L. 1991, ApJ, 366, 512

Cherepashchuk, A. M., Mihalas, D., & Pravdo, S. H. 2002, PASP, 114, 761

Cohen et al. 1997b; (8) Killian 1994; (9) Rho et al. 2001; (10) J. Rho, private communication; (11) Wojdowski et al. 2002; (12) Strickland 1995; (13) Waldron & Cassinelli 2001 (14) Miller et al. 2002.

#### TABLE 4

| Star | Spectral Type | Age Myr | Star Formation Region | $T$ (MK) | Magnetic Fields | $\log L_X^\text{wind}$ (ergs s$^{-1}$) | $\log L_X^\text{bol}$ (ergs s$^{-1}$) | $\log L_X^\text{wind}/L_X^\text{bol}$ | References |
|------|--------------|---------|-----------------------|--------|----------------|----------------------------------|---------------------------------|----------------------------------|------------|
| $\theta$ Ori A | B0.5 V | 0.3 | Orion | 5–43 | Yes | 31.0 | 30.7 | −7.3 | 1, 2, 3 |
| $\theta$ Ori B | B1 V/B3 | 0.3 | Orion | 22–35 | Probable | 30.3 | ? | ? | 2, 3, 4 |
| $\theta$ Ori C | O6.5Vp | 0.3 | Orion | 6–66 | Yes | 32.2 | 31.5 | −7.2 | 1, 2, 3 |
| $\theta$ Ori D | B0.5 Vp | 0.3 | Orion | 7–8 | ? | ? | 29.5 | −8.5 | 2, 3, 4 |
| $\theta$ Ori E | B0.5 | 0.3 | Orion | 4–47 | Yes | 31.4 | 30.8 | −7.2 | 1, 2, 3 |
| $\theta$ Ori | O9.5 Vpe | 0.3 | Orion | 5–32 | Yes | 31.1 | 31.4 | −7.1 | 5, 6 |
| $\gamma$ Sco | B0.2 V | 0.3 | Sco-Cen | 7–27 | Yes | 31.9 | 31.4 | 7 | 5, 6 |
| $\delta$ Ori | B0.5 | 0.3 | Orion | 5–47 | Yes | 31.4 | 31.7 | 7 | 5, 6 |
| HD 206267 | O6.5 V | 3–7 | IC 1396 | 2–10 | No | 31.6 | 37.2 | −11.2 | 11 |
| 15 Mon | O7 V | 3–7 | NGC 2264 | 3–12 | No | 32.3 | 32.3 | −7.2 | 1 |
| $\alpha$ Ori | O9 III | ≤12 | Orion | 1–10 | No | 31.7 | 31.7 | −7.2 | 1 |
| $\zeta$ Ori | O9.7 Ib | ≤12 | Orion | 1–10 | No | 32.4 | 32.4 | −6.8 | 1 |
| $\delta$ Ori | O9.5 II | ≤12 | Orion | 1–10 | No | 32.2 | 32.2 | −6.8 | 13, 14 |

The references for the table entries are as follows:

(1) This paper; (2) Hillenbrandt 1997; (3) Petr et al. 1998; (4) Schulz et al. 2001a; (5) Schulz et al. 2003; (6) Feigelson et al. 2002; (7) Cohen et al. 1997b; (8) Killian 1994; (9) Rho et al. 2001; (10) J. Rho, private communication; (11) Wojdowski et al. 2002; (12) Strickland 1995; (13) Waldron & Cassinelli 2001 (14) Miller et al. 2002.
Miller, N. A., Casinelli, J. P., Waldron, W. L., MacFarlane, J. J., & Cohen, D. H. 2002, ApJ, 577, 951
Nordsieck, K. H., Cassinelli, J. P, & Anderson, C. M. 1981, ApJ, 248, 678
O'dell, C. R., Wen, Z., & Hu, X. 1993, ApJ, 410, 696
Owocki, S. P., Castor, J. J., & Rybicki, G. B. 1988, ApJ, 335, 914
Owocki, S. P., & Cohen, D. H. 2001, ApJ, 559, 1108
Pallavicini, R., Golub, L., Rosner, R., Vaiana, G. S., Ayres, T., & Linsky, J. L. 1981, ApJ, 248, 279
Pauldrach, A. W., Kudritzki, R. P., Puls, J., Butler, K., & Hunsinger, J. 1994, A&A, 283, 525
Petr, M. G., Coudé du Foresto, V., Beckwith, S. V. W., Richichi, A., & McCaughrean, M. J. 1998, ApJ, 500, 825
Predehl, P., & Schmitt, J. H. M. M. 1995, A&A, 293, 889
Puls, J., Owocki, S. P., & Fullerton, A. W. 1993, A&A, 279, 457
Rho, J., Corcoran, M. F., Chu, Y.-H., & Reach, W. T. 2001, ApJ, 562, 446
Reimers, A., Stahl, O., Wolf, B., Kaufer, A., & Szeifert, T. 2000, A&A, 363, 585
Runacres, M. C., & Owocki, S. P. 2002, A&A, 381, 1015
Savage, B. D., & Jenkins, E. P. 1972, ApJ, 172, 491
Schmelz, J. T., Saba, J. L. R., & Strong, K. T. 1992, ApJ, 398, L115
Schulz, N. S. 2002, Rev. Mexicana Astron. Astrofis. Ser. Conf. 15, 220
Schulz, N. S., Canizares, C. R., Huenemoerder, D., Kastner, J. H., Taylor, S. C., & Bergstrom, E. J. 2001, ApJ, 549, 441
Schulz, N. S., et al. 2003, ApJ, submitted
Schulz, N. S., Canizares, C. R., Huenemoerder, D., & Lee, J. 2000, ApJ, 545, L135
Schulz, N. S., Huenemoerder, D., Kastner, J. H., & Lee, J. 2001a, BAAS, 33(4), 2205
Shu, F. H., Shang, H., Glassgold, A. E., & Lee, T. 1997, Science, 277, 1475
Seward, F. D, Froman, W. R., Giaconi, R., Griffith, R. E., Harnden, F. R., Jr., Jones, C., & Pye, J. P. 1979, ApJ, 234, L55
Smith, R. K., Brickhouse, N. S., Liedahl, D. A., & Raymond, J. C. 2001, ApJ, 556, L91
Snow, T. P., & Jenkins, E. B. 1977, ApJS, 33, 269
Spitzer, L. 1978, Physical Processes in the Interstellar Medium (New York: Wiley)
Stahl, O., Wolf, B., Gäng, T., Gummersbach, C., Kaufer, A., Kovács, J., Mandel, H., & Szeifert, T. 1993, A&A, 274, L29
Stahl, O., et al. 1996, A&A, 312, 539
Strickland, D. J. 1995, Observatory, 115, 180
Tsuboi, Y., Imanishi, K., Koyama, K., Grosso, N., & Montmerle, T. 2000, ApJ, 532, 1089
Ud-Doula, A., & Owocki, S. P. 2002, ApJ, 576, 413
Walborn, N., & Nichols, J. S. 1994, ApJ, 425, L29
Waldron, W. L., & Cassinelli, J. P. 2001, ApJ, 548, L45
Waljaski, K., Moses, D., Dere, K., Saba, J. L. R., Strong, K. T., Webb, D. F., & Zarro, D. M. 1994, ApJ, 429, 909
Weigelt, G., Balega, Y., Preibisch, T., Schertl, D., Schöller, M., & Zinnecker, H. 1999, A&A, 347, L15
Wojdowski, P. S., Schulz, N. S., Ishibashi, K., & Huenemoerder, D. 2002, in High-Resolution X-Ray Spectroscopy with XMM-Newton and Chandra, ed. G. Branduardi-Raymont (Univ. College London: Mullard Space Sci. Lab.)
Yamauchi, S., & Kamimura, R. 1999, Star Formation 1999, ed. T Nakamoto (Nobeyama: Nobeyama Radio Obs.), 308
Yamauchi, S., & Koyama, K. 1993, ApJ, 404, 620
Yamauchi, S., Koyama, K., Nakano, M., & Okada, K. 1996, PASJ, 48, 719