The corona and upper transition region of $\epsilon$ Eridani

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ABSTRACT

We present analyses of observations of $\epsilon$ Eridani (K2 V) made with the Low Energy Transmission Grating Spectrometer onboard Chandra and the Extreme Ultraviolet Explorer, supplemented by observations made with the Space Telescope Imaging Spectrograph, the Far Ultraviolet Spectroscopic Explorer and the Reflection Grating Spectrometer onboard XMM–Newton. The observed emission lines are used to find relative element abundances, to place limits on the electron densities and pressures and to determine the mean apparent emission measure distribution. As in a previous paper by Sim & Jordan, the mean emitting area as a function of the electron temperature is derived by comparisons with a theoretical emission measure distribution found from energy balance arguments. The final model has a coronal temperature of $3.4 \times 10^6$ K, an electron pressure of $1.3 \times 10^{16}$ cm$^{-3}$ K at $T_e = 2 \times 10^5$ K and an area filling factor of 0.14 at $3.2 \times 10^5$ K. We discuss a number of issues concerning the atomic data currently available. Our analyses are based mainly on the latest version of CHIANTI (v5.2). We conclude that the Ne/O relative abundance is 0.30, larger than that recommended from solar studies, and that there is no convincing evidence for enhanced coronal abundances of elements with low first ionization potentials.

Key words: stars: abundances – stars: coronae – stars: individual: $\epsilon$ Eridani – stars: late-type.

1 INTRODUCTION

$\epsilon$ Eridani ($\epsilon$ Eri) (K2 V) is a nearby dwarf star that has been observed over a wide spectral range, from the infrared to X-ray wavelengths. Its fundamental parameters have been discussed by Drake & Smith (1993), where references to earlier work can be found, and more recently by Di Folco et al. (2004) and Allende Prieto et al. (2004). Di Folco et al. (2004) have derived the diameter of $\epsilon$ Eri from interferometric measurements and have also made models to check the self-consistency of their adopted parameters. A more recent measurement by Di Folco et al. (2007) gives a radius that is smaller by only 0.008 $R_\odot$. The values of the parameters adopted here are given in Table 1 and are from Di Folco et al. (2004), although these are also consistent with those found by Drake & Smith (1993), to within the combined uncertainties. Compared with the Sun, $\epsilon$ Eri has a shorter rotational period (11.68 d) (Donahue, Saar & Baliunas 1996), a larger spatially averaged magnetic field (165 G) (Rüedi et al. 1997) and larger stellar surface emission-line fluxes. It is therefore ideal for studies of a stellar outer atmosphere under conditions of higher mean magnetic activity than the Sun.

Observations with IRAS showed that $\epsilon$ Eri has an infrared excess, indicating the presence of dust; more recently a planet and a debris disc have been detected (Hatzes et al. 2000; Greaves et al. 2005).

Further studies of the dust have been made by Di Folco et al. (2007). Combining the above rotational period with $\upsilon\,\sin\,i = 1.7$ km s$^{-1}$ and $R_\star = 0.743$ $R_\odot$ (Di Folco et al. 2004), leads to $i = 32^\circ$, similar to the value of 25$^\circ$ suggested by Greaves et al. (2005) for the inclination of the debris disc to the plane of the sky.

Here, we make use of observations at ultraviolet (UV) to X-ray wavelengths, as summarized in Table 2. In previous work, we used observations with the Space Telescope Imaging Spectrograph (STIS) onboard the Hubble Space Telescope to identify forbidden lines of Fe XII (Jordan et al. 2001a) and to determine the electron pressure ($P_e$) in the transition region (Jordan et al. 2001b). Including observations with the Far Ultraviolet Spectroscopic Explorer (FUSE), we have modelled the chromosphere and transition region, up to an electron temperature ($T_e$) of $\gtrsim 3 \times 10^5$ K (Sim & Jordan 2005).

Observations with the Extreme Ultraviolet Explorer (EUV) have been analysed by a number of authors. Schmitt et al. (1996) used lines of iron to derive the emission measure distribution (EMD) at above $T_e \simeq 6 \times 10^5$ K and also placed limits on the electron density ($n_e$) in the upper transition region from lines of Fe XIII and Fe XIV. Laming, Drake & Widing (1996) used the same spectra to derive emission measures from lines of a range of elements and concluded that if any trends in abundances with the first ionization potential (FIP) were present, they were not significantly larger than that in the solar corona. Sanz-Forcada, Brickhouse & Dupree (2003) included further EUVE spectra and also derived the EMD. They found much
2 OBSERVATIONS AND DATA REDUCTION

2.1 X-ray observations

We use the Chandra LETGS spectra (ObsID 1869) that had an exposure time of 105.3 ks. We have also examined the spectra obtained with the XMM–Newton RGS instrument (ObsID 0112880501) that have an exposure time of 13 ks. These spectra are available from the Chandra and XMM–Newton archives.

The LETGS projects the dispersed spectrum on to a microchannel plate detector. The detector is placed behind the grating in a manner such that the non-dispersed photons (zeroth-order) are recorded in the middle of the detector, with the two dispersion directions appearing as spectra in the plus and minus directions. There is overlap between the first-order spectra and the various higher order spectra of shorter wavelengths and care must be taken to exclude higher order lines or to account for line blending when this occurs in important first-order lines. In the lists of lines given in Tables 3 and 4, purely higher order lines are excluded and blends with first-order lines are noted.

2.2 Data reduction

We extracted the LETGS spectra on the plus and minus sides separately and calculated the effective areas using the standard CIAO tools (v3.2). The measurement of line fluxes was carried out with the cora program developed by Ness & Wichmann (2002). This accounts for the particular problems of low-count photon statistics (see e.g. Ness et al. 2001). After trying several approaches, we used a fixed linewidth (full width at half-maximum – FWHM) of 0.060 Å, since intrinsic stellar linewidths are not resolvable.

As noted by others (see the extensive discussion in Chung et al. 2004), the apparent wavelengths of lines can differ between the plus and minus sides, leading to larger linewidths when the spectra in the plus and minus sides are summed. The line fluxes given in Table 3 and about half of those given in Table 4 are derived from the summed spectra, using variable linewidths of between 0.053 and 0.080 Å, although the larger widths usually occur only at wavelengths above 90 Å. We have checked that these fluxes do not differ significantly from the averages of the fluxes in the individual plus- and minus-side spectra. Some lines are not observed, or are poorly observed, in one of the two spectra. In this case, only one spectrum is used; the wavelengths of such lines are given in italic script in Table 4. We also give the effective areas adopted so that the original count rates can be recovered.

All the line fluxes have then been corrected for absorption in the interstellar medium (ISM) using a hydrogen column density of log \( N_H (\text{cm}^{-2}) = 17.88 \) from Dring et al. (1997), lower than that used by Sanz-Forcada et al. (2003) who adopted a value of 18.1. The absorption model also includes He I, He II and other abundant elements. Uncertainties in the line fluxes arise not only from the statistical measurement errors calculated by the cora program, but also from the linewidths adopted and the source continuum. The linewidths in the LETGS spectra can be approximated by Moffat profiles (Lorentzians with an exponent \( \beta = 2.4 \); Drake 2004). The source continuum can be accounted for in cora by adding a background (in counts Å\(^{-1}\)) to the line templates. By varying the source background, we found that in many cases, the dominant source of the uncertainty in line fluxes arises from the choice of the source background. The specific case of the lines of O vii is discussed in Section 3.
The APED (Smith et al. 2001) and CHIANTI (v5.2) (Landi et al. 2006) data bases were consulted in making the line identifications (see also Section 4). Lines marked with b are possibly blended but the identification of the main additional contributor is not certain and these lines are not used in the analyses of line fluxes. Blending between first-order lines occurs in several important cases. The methods used to find the relative contributions are discussed in Section 4. Fluxes were measured for all lines, but since our aim is to establish a reliable EMD a number of weak lines are not included in Tables 3 and 4, unless they have particular significance for our

| λ (Å) | Flux | A_{eff} (cm^2) | λ" (Å) | Ion | Transition |
|-------|------|----------------|-------|-----|------------|
| 6.65  | 8.39 ± 2.77 | 44.1 | 6.65 | Si xiii | 1s^2 1S0–1s 2p 1P1 |
| −6.74 | 6.69 | Si xiii | 1s^2 1S0–1s 2p 1P1 |
| 8.42  | 3.68 ± 0.66 | 38.2 | 8.42 | Mg xii | 1s^2 1S0–2p 2P1/2 |
| 9.17  | 12.6 ± 2.5 | 32.5 | 9.17 | Mg xii | 1s^2 1S0–1s 2p 1P1 |
| −9.31 | 9.23 | Mg xii | 1s^2 1S0–1s 2p 1P1 |
| 10.23 | 4.20 ± 0.75 | 28.7 | 10.24 | Ne x | 1s^2 1S0–2p 1P1 |
| 11.27 | 4.69 ± 0.74 | 28.5 | 11.26 | Fe xvii | 2p^6 1S0–2p 2P3/2 D3/2 |
| 11.55 | 5.58 ± 0.75 | 29.1 | 11.55 | Ne ix | 1s^2 1S0–1s 2p 1P1 |
| 12.14 | 29.5 ± 1.4 | 28.7 | 12.14 | Ne x | 1s^2 1S0–2p 1P1 |
| 12.29 | 7.18 ± 0.80 | 28.5 | 12.26 | Fe xvii | 2p^6 1S0–2p 2P3/2 D1 |
| 13.45 | 22.9 ± 1.05 | 29.4 | 13.45 | Ne ix | 1s^2 1S0–1s 2p 1P1 |
| 13.55 | 7.57 ± 0.60 | 29.4 | 13.55 | Ne ix | 1s^2 1S0–1s 2p 1P1 |
| 13.70 | 16.1 ± 0.90 | 29.4 | 13.70 | Ne ix | 1s^2 1S0–1s 2p 1P1 |
| 13.83 | 3.67 ± 0.41 | 29.4 | 13.82 | Fe xvii | 2s^2 2p^6 1S0–2s 2p 3P0/2 |
| 14.21 | 10.9 ± 0.9 | 29.5 | 14.21 | Fe xviii | 2p^3 3P0–2p 2D3/2 |
| −14.26 | 14.26 | Fe xviii | 2p^3 3P0–2p 2D3/2 |
| 14.38 | 4.19 ± 0.69 | 29.6 | 14.37 | Fe xviii | 2p^3 3P0–2p 2D3/2 |
| 15.02 | 52.1 ± 1.6 | 30.3 | 15.02 | Fe xvii | 2p^6 1S0–2p^5 3P0 3P1 |
| 15.19 | 7.98 ± 0.84 | 30.4 | 15.18 | O viii | 1s^2 2s1–4p 2P1/2 |
| 15.27 | 21.8 ± 1.1 | 30.5 | 15.26 | Fe xvii | 2p^6 1S0–2p^5 3P0 3P1 |
| 16.01 | 1.7 ± 1.0 | 29.9 | 16.01 | O viii | 1s^2 3P0–2p 2D3/2 |
| 16.09 | 5.80 ± 0.67 | 30.2 | 16.08 | Fe xviii | 2p^2 3P0–2p 2P1/2 |
| 16.77 | 30.4 ± 1.2 | 30.6 | 16.78 | Fe xvii | 2p^2 3P0–2p 2P1/2 |
| 17.05 | 77.0 ± 3.0 | 25.7 | 17.05 | Fe xvii | 2p^2 3P0–2p 2P1/2 |
| 17.10 | 17.10 | Fe xvii | 2p^2 3P0–2p 2P1/2 |
| 18.63 | 7.79 ± 1.26 | 26.3 | 18.63 | O vii | 1s^2 1S0–1s 2p 1P1 |
| 18.97 | 88.2 ± 1.9 | 26.5 | 18.97 | O viii | 1s^2 1S0–2p 2P1/2 |
| 21.61 | 41.5 ± 1.5 | 17.3 | 21.60 | O vii | 1s^2 1S0–1s 2p 1P1 |
| 21.81 | 9.60 ± 0.83 | 17.0 | 21.81 | O vii | 1s^2 1S0–1s 2p 1P1 |
| 22.21 | 27.3 ± 1.3 | 17.0 | 22.20 | O vii | 1s^2 1S0–1s 2p 1P1 |
| 24.79 | 10.8 ± 0.9 | 16.8 | 24.78 | O vii | 1s^2 1S0–1s 2p 1P1 |
| 28.47 | 3.53 ± 0.55 | 15.6 | 28.47 | C vii | 1s^2 1S0–3P1/2 |
| 28.79 | 3.56 ± 0.55 | 15.3 | 28.79 | N vii | 1s^2 1S0–3P1/2 |
| 29.09 | <2.50 | 15.0 | 29.08 | N vii | 1s^2 1S0–3P1/2 |
| 29.54 | <2.50 | 14.1 | 29.53 | N vii | 1s^2 1S0–3P1/2 |
| 30.45 | 4.37 ± 0.64 | 12.1 | 30.47 | S xiv | 2s^2 3S1–2P1/2 |
| 32.24 | 2.47 ± 0.52 | 13.5 | 32.24 | S xiii | 2s^2 3S1–2P1/2 |
| 32.56 | 3.70 ± 0.55 | 13.3 | 32.56 | S xiv | 2p^2 3P0–3S0 3S1 |
| 33.55 | 2.47 ± 0.49 | 12.8 | 33.55 | S xiv | 2p^2 3P0–3S0 3S1 |
| 33.74 | 17.1 ± 1.0 | 12.8 | 33.74 | C vii | 1s^2 1S0–2p 2P1/2 |
| 35.69 | 4.17 ± 0.55 | 12.3 | 35.67 | S xiii | 2s^2 3P0–3s 3P0 |
| 37.61 | 2.14 ± 0.49 | 10.4 | 37.60 | S xiv | 2p^2 3P0–3S0 3S1 |
| 40.27 | 3.35 ± 0.78 | 5.5 | 40.27 | C vii | 1s^2 1S0–1s 2p 1P1 |
| 43.76 | 2.89 ± 0.27 | 25.9 | 43.76 | Si xii | 2s^2 3S1–2p 3P0 |
| 44.17 | 3.75 ± 0.29 | 26.1 | 44.02 | Si xii | 2p^2 3P0–3S0 3S1 |
| 45.51 | 1.30 ± 0.21 | 25.7 | 45.52 | Si xii | 2p^2 3P0–3S0 3S1 |
| 45.69 | 2.01 ± 0.24 | 25.6 | 45.69 | Si xii | 2p^2 3P0–3S0 3S1 |

*Fluxes are in 10^{-14} erg cm^{-2} s^{-1}. bLikely to be blended. cThe percentage contribution to the total flux given. dFluxes after deblending.

Wavelengths from CHIANTI (v5.2).
Table 4. Continuation from Table 3.

| λ (Å) | Flux* | A_{eff} (cm^2) | λ' (Å) | Ion | Transition |
|------|-------|---------------|--------|-----|------------|
| 49.21 | 3.79 ± 0.29 | 24.4 | 49.22 | Si xi | 2s 2p^1P_1–2s 3d^1D_2 |
| 50.36 | 4.96 ± 0.49 | 10.9 | 50.36 | Fe xvi | 3s^2S_{1/2}–4p^2P_{3/2} |
| ~29 per cent | | | 50.34 | | 3rd order 16.78 Å |
| 50.55 | 3.75 ± 0.44 | 10.9 | 50.52 | Si x | 2p^2P_{1/2}–2s^3D_{3/2} |
| ~51 per cent | | | 50.57 | Fe xvi | 3s^2S_{1/2}–4p^2P_{1/2} |
| 50.69 | 2.83 ± 0.39 | 10.9 | 50.69 | Si x | 2p^2P_{1/2}–2s^3D_{3/2} |
| 52.32 | 2.16 ± 0.36 | 10.4 | 52.30 | Si xi | 2s 2p^1P_1–2s 3s^1S_0 |
| 52.90 | 2.24 ± 0.37 | 10.4 | 52.91 | Fe xv | 3s^2S_{1/2}–3s 4p^1P_1 |
| 54.14 | 2.84 ± 0.38 | 10.2 | 54.13 | Fe xvi | 3p^2P_{1/2}–4d^3D_{3/2} |
| 54.72 | 4.78 ± 0.46 | 10.1 | 54.71 | Fe xvi | 3p^2P_{3/2}–4d^3D_{3/2} |
| 57.90 | 2.11 ± 0.48 | 9.2 | 57.92 | Mg x | 2s^2S_{1/2}–3p^2P_{1/2} |
| 59.40 | 2.97 ± 0.44 | 8.0 | 59.40 | Fe xvi | 3s^2D_{3/2}–3s 4f^1D_2 |
| 62.88 | 2.73 ± 0.43 | 7.7 | 62.87 | Fe xvi | 3p^2P_{1/2}–4s^3S_{1/2} |
| 63.15 | 1.07 ± 0.34 | 7.7 | 63.15 | Mg x | 2p^2P_{1/2}–2s^3D_{3/2} |
| 63.31 | 2.80 ± 0.44 | 7.7 | 63.31 | Fe xvi | 3p^2P_{1/2}–2s^3D_{3/2} |
| 63.73 | 6.20 ± 0.57 | 7.6 | 63.71 | Fe xvi | 2p^2P_{3/2}–4s^3S_{1/2} |
| 65.86 | 1.06 ± 0.37 | 7.3 | 65.85 | Mg x | 2p^2P_{1/2}–2s^3S_{1/2} |
| 66.25 | 3.78 ± 0.51 | 7.2 | 66.25 | Fe xvi | 3d^2D_{3/2}–2s^3F_{5/2} |
| 66.35 | 5.91 ± 0.60 | 7.1 | 66.36 | Fe xvi | 3d^2D_{3/2}–4f^2P_{3/2} |
| 69.60 | 1.07 ± 0.25 | 14.3 | | | b |
| 69.68 | 5.47 ± 0.41 | 14.3 | 69.68 | Fe xv | 3s 3p^1P_1–3s 4s^1S_0 |
| 72.32 | <1.14 | 13.5 | 72.31 | Mg ix | 2s 2p^1P_1–2s 3d^1D_2 |
| 73.48 | 2.70 ± 0.55 | 13.0 | 73.47 | Fe xv | 3s 3d^1D_2–3s 4f^1F_3 |
| 76.04 | 1.32 ± 0.27 | 12.2 | 76.02 | Ne viii | 2p^2P_{1/2}–2s^3D_{3/2} |
| 76.13 | 1.21 ± 0.27 | 12.2 | 76.15 | Fe xiv | 3d^2D_{3/2}–2s^3F_{5/2} |
| 76.53 | 1.61 ± 0.28 | 12.0 | | | ? |
| 77.74 | <0.63 | 11.7 | 77.74 | Fe xvi | 3d^2D_{3/2}–4p^2P_{1/2} |
| 88.11 | 5.07 ± 0.89 | 9.5 | 88.08 | Mg ix | 2s^2P_{1/2}–2s 3s^1S_0 |
| 88.15 | 5.07 ± 0.89 | 9.5 | 88.08 | Ne viii | 2s^2S_{1/2}–2p^2P_{3/2} |
| 93.97 | 7.08 ± 0.51 | 8.8 | 93.92 | Fe xvii | 2s^2p^2P_{3/2}–2s 2p^2P_{3/2} |
| 98.15 | 1.68 ± 0.34 | 7.6 | 98.12 | Ne viii | 2p^2P_{3/2}–3d^3D_{3/2} |
| 98.28 | 3.90 ± 0.45 | 7.6 | 98.26 | Ne viii | 2p^2P_{3/2}–3d^3D_{3/2} |
| 101.62 | 1.43 ± 0.36 | 7.1 | 101.55 | Fe xix | 2s^2p^2P_{3/2}–2s 2p^2P_{1/2} |
| 104.00 | 2.57 ± 0.41 | 6.9 | 103.94 | Fe xviii | 2s^2p^2P_{1/2}–2s 2p^2P_{1/2} |
| 108.42 | 2.54 ± 0.41 | 6.8 | 108.36 | Fe xix | 2s^2p^2P_{3/2}–2s 2p^2P_{1/2} |
| 132.93 | 3.9 | 132.84 | Fe xx | 2s^2p^2P_{3/2}–2s 2p^2P_{1/2} |
| 141.09 | 2.57 ± 0.57 | 4.0 | 141.04 | Ca xii | 2s^2p^2P_{3/2}–2s 2p^2P_{1/2} |
| 150.16 | <2.17 | 3.6 | 150.12 | O vi | 2s^2S_{1/2}–2p^2P_{1/2} |
| 152.25 | 2.82 ± 0.57 | 3.6 | 152.15 | Ni xii | 3p^2P_{3/2}–3p^2P_{3/2} |
| 154.25 | 1.57 ± 0.50 | 3.6 | 154.16 | Ni xii | 3p^2P_{3/2}–3p^2P_{3/2} |
| 171.17 | 13.0 ± 1.7 | 1.2 | 171.07 | Fe x | 3s^6S_{0/2}–3p^3d^1P_1 |

*Fluxes are in 10^{-14} erg cm^{-2} s^{-1}. b Likely to be blended. The percentage contribution to the total flux given. d Wavelengths in italics indicate flux measurements on one side only. c Wavelengths from CHIANTI (v5.2).

studies. A fuller list of lines present has been published by Sanz-Forcada et al. (2004).

There are systematic differences between laboratory and observed wavelengths above about 80 Å (where the observed wavelengths are too large), owing to the treatment of detector-plate gaps (see Chung et al., 2004). Since we are not analysing line-shifts and are using only well-identified lines in this region, we have not attempted to correct these wavelengths.

In reducing the RGS1 and RGS2 spectra, we used Lorentzian line-profiles to approximate the instrumental line-profiles and used a fixed FWHM of 0.06 Å. The only use we have made of these spectra is in comparisons between LETGS and RGS fluxes where line-blending is not significant, because the instrumental wings to the lines prevent accurate deblending.

### 2.3 Other observations; potential variability of e Eri

In later sections, we will be using spectra obtained with the EUVE, STIS and FUSE. Table 2 gives the dates on which these observations were made.

Balunas et al. (1983, 1995) have studied the Ca II emission lines over various time-scales and find no clear single activity cycle.

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The monthly variations in the S-index cover a total amplitude of \( \pm 24 \) per cent, so that the chromospheric emission does not show substantial variations. Sim & Jordan (2005) found no significant differences between the fluxes in transition region lines observed with the International Ultraviolet Explorer (IUE) in 1981 and with the STIS in 2000.

Regarding variations in emission from the upper transition region and corona, Sanz-Forcada et al. (2003), using the EUVE Deep Survey Imager, observed similar mean levels of counts in the 80–180 Å passband in 1993 and 1995, although some variations of the order of \( \pm 30 \) per cent were observed over a time-interval of hours. Similarly, Sanz-Forcada et al. (2004) show light curves obtained from the LETGS observations that exhibit variations amounting to about \( \pm 21 \) per cent over 28 h. The only systematic difference between the LETGS and EUVE observations is that the fluxes of lines of iron in stages of ionization greater than Fe XVIII are lower in the LETGS spectra, suggesting that the star had less hot active region material at that time. For the lines observed in both the RGS1 and RGS2 spectra, the mean RGS fluxes are on average a factor of only 1.06 larger than those derived from the LETGS spectra. We conclude that significant uncertainties should not be introduced by combining observations made on different dates.

### 3 THE ELECTRON PRESSURE

Jordan et al. (2001b) measured a number of density-sensitive line flux ratios in their STIS spectrum of \( \epsilon \) Eri and used these to investigate \( P_e \), and to assess the available atomic data (in CHIANTI v3.01 and relevant papers). Throughout this work, we define \( P_e = N_e T_e \) cm\(^{-3}\) K. Values of \( \log P_e = 15.69 \pm 0.1 \) at \( T_e = 4.5 \) and \( \pm 15.67 \) at \( T_e = 5.1 \) were obtained from the transition region lines of Si III and O IV, respectively, although for O IV, the four independent ratios gave values of \( \log P_e \) between \( 15.28 \pm 0.08 \) and \( 16.16 \pm 0.17 \). The overall mean value of \( \log P_e = 15.68 \) was consistent with the lower and upper limits provided by lines of C iii, O v and Fe xv.

We have re-examined all the pressures derived using CHIANTI (v5.2), since some changes were made following the discussions in Jordan et al. (2001b). There have been no changes to the data for the lines of C iii and Si iii. The pressures found from the ratios involving the O IV 1401-Å line are now slightly lower (\( \log P_e = 15.28 \pm 0.08 \) and \( 16.16 \pm 0.17 \)), but the discordant low pressures derived from ratios involving the blended line at 1404 Å are unchanged, and there are still problems with the lines of Si iv. A pressure can now be found from the lines of O v at 1218 and 1371 Å, and log \( P_e = 15.95 \pm 0.15 \). If the pressures from the ratios involving the O IV 1401-Å line are preferred, then the mean electron pressure at around \( T_e = 5.3 \) becomes \( \pm 15.97 \pm 0.2 \). It should be noted that if material pressures up to about \( \log P_e = 16.30 \) were present, this could be detected from the line flux ratios used.

In models of the chromosphere and transition region by Sim & Jordan (2005), a turbulent pressure term derived from the observed non-thermal linewidths (Jordan et al. 2001b) is included in the equation of hydrostatic equilibrium. This results in small increases in \( P_e \) with increasing \( T_e \) above \( 3 \times 10^4 \) K, where the Si iii lines are formed. The pressures now derived from the transition region lines are consistent with this behaviour, to within the uncertainties given above. In our theoretical models of the upper transition region and inner corona (and in those by Sim & Jordan 2003a), \( P_e \) continues to rise by a few per cent until \( T_e \approx 5.8 \) (see Section 6). Only a small difference in \( P_e \) is expected between \( T_e = 5.3 \) and 6.3, where the lines of O vii are mainly formed.

In the X-ray spectra, the ratio of the fluxes in the forbidden line (f) \((1s 2s 3S–1s 2S 3S)\) and the intersystem line (i) \((1s 2p 3P_{1,2}–1s 2S 3S)\) in the He i-like ions is potentially sensitive to \( n_e \) (Gabriel & Jordan 1969). Of the He i-like ions observed with the LETGS, only O vii has lines that are sufficiently strong and unblended to use in a measurement of \( n_e \). Because of the importance of the measured flux ratio, we have varied the continuum level and linewidths used in extracting the fluxes, in order to obtain realistic error bars. The spectrum and these fits are shown in Fig. 1. The mean observed ratio is \( 2.88_{-0.34}^{+0.57} \). Provided that the same widths are used for both lines, uncertainties in the widths have less effect than those in the choice of continuum. The range of the observed ratio is \( 2.535–3.455 \), compatible with the ratio of 3.06 measured by Sanz-Forcada et al. (2004), who allowed explicitly for contributions from what they regarded as weak unidentified lines.

Using CHIANTI (v4.2), at \( T_e = 2 \times 10^4 \) K, the observed ratio of 2.88 leads to \( \log P_e = 16.35 \), with a range of 16.54–16.00. The lower limit is consistent with the pressure found from the transition region lines at \( \log T_e = 5.3 \). This pressure is essentially the same as the value of \( \log P_e = 16.33 \) found in an earlier analysis by Ness et al. (2002), which was adopted by Sanz-Forcada et al. (2004). (Using atomic data from APED gives essentially the same results.) CHIANTI (v4.2) predicts a ratio of 3.9 in the low-density limit, but does not include recombination to the \( n = 2 \) levels, either directly or via cascades. Early work by Gabriel & Jordan (1973) showed that including both radiative and di-electronic recombination tends to increase the predicted ratio in the low-density limit, but by only a relatively small amount; collisional excitation followed by cascades is more important. They predicted a low-density limit of 3.64. Blumenthal, Drake & Tucker (1972) found larger effects from recombination, but according to Gabriel & Jordan (1973) they overestimated the contributions from di-electronic recombination.

The most recent version of CHIANTI (v5.2) gives the option of including radiative recombination as a process populating the excited states. The implementation of radiative recombination in CHIANTI (v5.2) is not correct, since it does not take into account recombination to the \( 1s 2p 3P_{1,2} \) level (Landi et al. 2006) and hence omits an important process for populating the \( 1s 2s 3S \) level. This leads to
lower values of the f/i ratio at a given value of \( n_e \). In particular, the value of f/i at low densities (\( R_o \)) becomes \( \approx 3.35 \).

Porquet & Dubau (2000) have given radiative and di-electronic recombination rate coefficients and effective collision strengths for populating the \( n = 2 \) levels, including the effects of cascades from \( n > 2 \) in all cases. At log \( T_e = 6.3 \), their calculations predict a value of \( R_o = 3.82 \), a little smaller than that found from CHIANTI (v4.2). Compared with the work by Gabriel & Jordan (1973), Porquet & Dubau (2000) find larger contributions from collisional excitations, followed by cascades, to levels with \( n > 2 \). Using the calculations by Porquet & Dubau (2000) (but neglecting the small contribution from recombination, since this causes only a small increase in \( R_o \)) leads to pressures that are similar to those from CHIANTI (v4.2) (log \( P_e = 16.38 \), with a range from 16.56 to 15.93). Here, the lower limit is compatible with the pressure found at around log \( T_e = 5.3 \), without considering the error bars.

At present, the origin of the higher optimum pressure found from the O VII lines is not clear, but we think that it is in part due to remaining uncertainties in the atomic data, as well as those in the flux measurements. A fuller atomic model for the He I-like ions is clearly needed in CHIANTI. Ideally, the value of \( R_o \) should be established from observations of the quiet solar corona, where the density is expected to be sufficiently low to give this limiting ratio. Gabriel & Jordan (1973) used an observed ratio of 3.6 in this manner, although Blumenthal et al. (1972) quote higher ratios of 3.78 and 3.92. Unfortunately, the LETGS spectra of α Cen A (Ness et al. 2002), where a solar-like pressure might be expected, do not have sufficient flux in the O VII lines to measure \( R_o \) to within useful limits.

The EUVE lines of Fe XIV are formed around log \( T_e = 6.25 \), within the range over which the O VII lines are formed. Laming et al. (1996) found log \( P_e = 15.25 \) from the lines at 211.33 and 219 Å, but the latter is weak and blended. Using the 211.33- and 246.78-Å lines, Schmitt et al. (1996) found log \( P_e = 15.55 \), still lower than that indicated by the O VII and transition region lines.

A number of other line ratios are sensitive to \( n_e \), because the relative populations of the levels in the ground term are not given by the Boltzmann population (e.g. lines in the B I-like isoelectronic sequence and lines of iron). The densities derived depend on the overall form of the EMD and relative element abundances and are discussed in Sections 4.5 and 4.6.1.

In the calculations that follow in Sections 4 and 5, we have explored the results using pressures of log \( P_e = 15.30, 15.68 \) and 16.10. In Section 6, we require the theoretical models to produce a value of log \( P_e = 15.97 \pm 0.20 \) at log \( T_e = 5.3 \).

### 4 Emission Measure Loci and Line Identifications

The identifications of the strong lines in the LETGS range are well known. In identifying other lines, and to check for blends, we used both APED (ATOMDB v1.3.1) and the CHIANTI data base (v5.2) (Landi et al. 2006) to explore which transitions might be present at a given wavelength. Emission measure loci (EMLs) were then calculated for possible candidates, including any dependence on \( P_e \).

For a spherically symmetric atmosphere, the line flux observed at the Earth is given by

\[
F_{21} = \frac{R_e^2 \cdot h \cdot n_E}{d^2 \cdot \lambda} \int n_2 \cdot A_{21} \cdot \frac{n_{\alpha}}{n_e} \cdot G(r) \cdot f(r) \cdot n_i \cdot n_H \cdot d\rho,
\]

where \( G(r) \) is the fraction of photons not intercepted by the star, \( f(r) = r^2 / R_e^2 \) and \( d \) is the distance to the star. The excited and lower levels are 2 and 1, respectively (where 1 is not necessarily the ground state); \( n_2 / n_1 \) is the abundance of the element relative to hydrogen, taken as constant over the region of line formation; \( n_{\alpha} / n_E \) is the relative ion population; \( n_H \) is the hydrogen number density; \( A_{21} \) is the spontaneous transition probability and the integration is over the radial distance, \( d \).

Equation (1) can be rewritten as

\[
F_{21} \approx g(n_e, T_e) \int n_2 n_0 G(r) f(r) \cdot d\rho,
\]

where \( g(n_e, T_e) \) includes all other terms in equation (1). Provided any dependence on \( P_e \) is taken into account, the EML gives the value of the apparent emission measure \( |EM| = \int n_\alpha n_H G(r) f(r) \cdot d\rho \) that would be required to account for all the observed flux, at each value of \( T_e \) in turn. The loci therefore provide useful constraints on the mean EMD, since if this exceeds the minimum of a locus by more than a small factor, too much flux will be predicted when the mean EMD is used to predict the line fluxes.

It is important to note that loci from lines of different isoelectronic sequences can have different variations of \( g(T_e) \) with \( T_e \), owing to the systematic differences in \( n_{\alpha} / n_E \) as a function of \( T_e \). Otherwise, if a line with a broad \( g(T_e) \) function were compared with one with a narrow \( g(T_e) \) function, an incorrect relative element abundance would be deduced. These differences are taken into account in finding the mean EMD (see Section 5).

Values of \( n_2 A_{21} / n_E \) have been calculated using CHIANTI (v5.2). The relative ion populations for iron have been taken from Arnaud & Raymond (1992). For O VII, the calculations given in Sim & Jordan (2005) have been adopted, since these include the density dependence of di-electronic recombination. All other values are taken from Arnaud & Rothenflug (1985). The element abundances initially adopted and the corrections to these required by the observations are discussed in Section 5.

The lines of elements other than iron are now discussed according to their isoelectronic sequence. Comparisons have been made between observed and predicted line flux ratios (or relative EMLs) using a single temperature for the line formation and also using the total fluxes predicted using the final EMD. Unless otherwise stated, both approaches give the same results. Although only the lines that we regard as the most reliable are used to determine the mean EMD, we calculate the predicted fluxes in all the lines discussed below and later compare these with the observed fluxes.

#### 4.1 Hydrogen-like lines

The (unresolved) Lyman α lines of C VII, N VII, O VIII, Ne X and Mg XII are all observed, although the Mg XII lines are weak. The Ne X lines at 12.13 + 12.14 Å are blended with a line of Fe XVII at 12.12 Å. Another line of Fe XVII is observed at 12.27 Å. The ratio of these two Fe XVII lines does not depend significantly on \( T_e \) so we have used the line at 12.27 Å to predict the flux in the line at 12.12 Å. This results in 73 per cent of the observed total flux in the line at 12.13 Å being due to Ne X.

The Lyman β lines are observed in O VIII (at 16.01 Å) and Ne X (at 10.23 Å), but the former is blended with a line of Fe XVIII at 16.00 Å. Another line of Fe XVIII at 16.07 Å has been used to find the contribution of the 16.00-Å line to the total flux. The ratio of the Fe XVIII lines is slightly sensitive to \( T_e \), and their temperature of optimum formation could lie between log \( T_e = 6.5 \) and 6.8. The total predicted fluxes show that 81 per cent of the observed line at 16.01 Å is due to O VIII. The ratio of the observed flux in the O VIII Lyman β line to that in the Lyman α line is then a factor of 1.2 larger than

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that predicted using the mean EMD. In Ne x, the observed Lyman \( \beta \)-to-Lyman \( \alpha \) ratio is a factor of 1.4 larger than predicted.

### 4.2 Helium-like lines

The resonance (r), intersystem (i) and forbidden line (f) of C v lie in a region of low and rapidly varying effective area and a real signal is observed only in the r line at 40.27 Å. However, this is slightly blended with the third order of the Ne ix r line and it is hard to determine a reliable flux (Ness et al. 2001). The uncertainty in the flux given in Table 3 includes the results of varying the background and linewidth. The r line of Nvi is present, although rather weak; the i and f lines are not sufficiently above the local noise level to be useful. The lines of O vii are relatively strong and their use in constraining \( P_n \) has been discussed in Section 3. Note that the contribution from radiative recombination has been included when predicting the fluxes in the singlet lines of all the He i-like ions.

The lines of Ne ix are also relatively strong, but are blended to various degrees with lines of Fe xvii, xviii and xix. A comparison between the LETGS spectra of \( \epsilon \) Eri and Capella, shown in Fig. 2, indicates that the blends with lines of iron are less important than in Capella. In \( \epsilon \) Eri, the Fe xv line at 13.83 Å is significantly weaker than the Ne ix f line and the Fe xix + xxi blend at 13.51 Å is significantly weaker than the r line. The deblending problem is therefore less severe than in Capella. The procedure used is as follows: the amplitudes and wavelengths of the lines observed in Capella using the High Energy Transmission Grating Spectrograph (HETGS) (see Ness et al. 2003a) are used to predict the counts in the LETGS spectra, taking into account the resolution and effective area of the LETGS; the lines of Fe xvii and Fe xviii are grouped together, as are the lines of Fe xix, while the lines of Ne ix are treated individually; and the amplitudes of the two groups and three individual lines are scaled iteratively to obtain the best fit to the LETGS spectrum of \( \epsilon \) Eri and hence the fluxes in the Ne ix lines are determined. The fits made are shown in Fig. 2. The analyses of all lines of neon and other lines of iron show that the relative weakness of the blended iron lines in \( \epsilon \) Eri arises from both the lower EMD above \( \log T_e = 6.6 \) and a lower Fe/Ne relative abundance. Given the deblending process, in particular for the i line, it is difficult to make a reliable interpretation of the relative fluxes in the Ne ix r, i and f lines.

The r and f lines are present in Mg xi and Si xiii, but the i lines cannot be distinguished. In extracting fluxes, we have summed over the r, i and f lines. The predicted fluxes are lower limits, since the contribution to the populations of excited states from dielectronic recombination is not yet included in CHANTI (v5.2), and contributions from satellite lines are not yet included for Mg xi.

The 1s3p \(^1\)P–1s \(^1\)S transition in O vii is present as a weak line at 18.63 Å, but is blended with a weaker line. The same transition in Ne ix might also be present at 11.54 Å. These weak lines are not used in deriving the mean EMD.

### 4.3 Lithium-like lines

In the region below 170 Å the transitions observed in the Li i-like ions are those between the \( n = 3 \) and 2 levels (3p–2s, 3d–2p, 3s–2p). The 3p–2s (unresolved) lines in O vi at 150.1 Å are present in the minus-direction spectrum, but are barely above the instrumental noise level in the plus-direction spectrum. The measured flux is very sensitive to the background adopted and provides only an upper limit to the EML. Although the 3p–2s lines are not included in the derivation of the mean EMD, this reproduces the observed flux within a factor of 1.1. The 2p–2s transitions are both observed (around 1032 and 1038 Å) in spectra obtained with the FUSE in 2000 December. They are useful in constraining the EMD between \( \log T_e = 5.5 \) and 6.1 (see Section 5).

In Ne viii, the 3p–2s and 3d–2p transitions are observed at around 88.1 and 98.2 Å, respectively. Although the 3s–2p transitions are just present around 103.1 Å, they are too weak to yield a reliable flux. The 4d–2p transitions occur around 73.5 Å but on the basis of the calculated EMLs, the line observed is identified as one of Fe xv. Heroux, Cohen & Malinovsky (1972) suggested that in the solar spectrum, the lines around 88.1 Å are blended with lines of Fe xi and identify a line at 86.88 Å with the strongest member of the multiplet (at solar densities). At the higher value of \( n_e \) in \( \epsilon \) Eri, the line at 86.88 Å should still be the strongest member of the multiplet, but is not observed. We have rejected the possibility of a second-order line of Si xii at 88.04 Å because a stronger second-order Si xii line at 88.34 Å is not present. Thus, significant blending with these Ne viii lines seems unlikely. Note that the relevant lines of Fe xi are not yet included in either CHANTI or APED. Heroux et al. (1972) also suggested that both the lines at 98.11 and 98.26 Å are blended with other lines in the solar spectrum. In \( \epsilon \) Eri, to within the measurement uncertainties, the two lines have the expected wavelength separation and almost the theoretical flux ratio, so there is no obvious evidence of substantial blending.

The lines in Mg x are all weak. The 3p–2s multiplet at 57.90 Å is not resolved; only the summed flux is used. The 3d–2p transitions at 63.30 + 63.31 and 63.16 Å are observed only in the minus-direction spectrum, owing to a chip gap in the plus direction. We use only the blend of transitions at around 63.3 Å in deriving the EMD, since the weaker line at 63.16 Å is barely above the noise level. The 3s–2p lines at 65.67 and 65.85 Å have an incorrect flux ratio in the minus-side spectrum and summed spectra, suggesting that the intrinsically weaker line at 65.67 Å is blended. Also, the line at 65.85 Å might be blended with another line at around 65.93 Å making it difficult to extract a reliable flux; these two Mg x lines are excluded from the derivation of the mean EMD.

In Si xii, the 3p–2s lines lie around 40.93 Å in the region of the instrumental absorption edge, and are not observable. The 3d–2p transitions at 44.02 and 44.17 Å are observed as moderately strong
lines. The observed ratio of the stronger to weaker components is about 1.5, instead of the predicted value of 1.99. We use only the stronger line, on the grounds that the weaker line might be blended. The 3s–2p lines at 45.52 and 45.69 Å are also observed, but are not used in deriving the mean EMD, as they are weak and blended with other weak lines or instrumental noise.

There are weak lines that are possibly due to S XIV. A line at 32.56 Å corresponds to the stronger component of the 3d–2p transitions. The 2p3d /2–2p 3/2 transition in Ca XII occurs at 30.47 Å and could contribute to the measured flux. These lines were not used in deriving the mean EMD.

In Ne VIII, the ratio of the observed flux from the 3p–2s transitions to that in the 3d–2p transitions is about a factor of 1.5 larger than predicted. This suggests that the atomic models or atomic data for the lithium-like ions would bear closer examination.

4.4 Beryllium-like lines

No lines of Mg IX are observed. Because it is useful to constrain the EMD at around 10^4 K, we have used the background level at 72.31 and 77.74 Å (the wavelengths of the 2s3d–2p and 2s3s–2s2p transitions, respectively), to find the upper limit to their EMLs.

Three singlet lines of Si X are observed; the 2s3p–2s2 line at 43.76 Å, the 2s3d–2p line at 49.22 Å and the 2s3s–2s2p line at 52.30 Å (but the latter only in the plus-side spectrum). Only the 52.30-Å line appears to be unblended.

The strongest transition (2s 3d–2s2p) in S XII is observed at 35.67 Å; other transitions are present, but are barely above the noise level.

4.5 Boron-like lines

The 3d–2p lines of Si X lie at 50.69 Å (2P3/2–2D5/2, 3/2) and 50.52 Å (2P1/2–2D3/2). As in Fe XIV (Jordan 1965), the ratio of these lines depends on n_e, since the relative populations of the ground 1sP levels have a Boltzmann distribution only at large values of n_e. The line at 50.52 Å is blended with a line of Fe XVI at 50.55 Å and careful deblending is required. Using the procedure described in Section 4.6.3, the ratio of the deblended fluxes in the 50.69- and 50.52-Å lines is 1.52 ± 0.26. Using CHIANTI (v5.2) and the mean EMD, the predicted integrated fluxes give ratios of 1.51 at log P_e = 15.68, 1.18 at log P_e = 15.30 and 1.78 at log P_e = 16.10. Thus, a value of log P_e close to 15.68 gives the best fits, but the uncertainty in the flux ratio just includes log P_e = 16.10. The flux in the unblended Si X line at 50.69 Å is also weakly dependent on n_e (compared with the n_e dependence of most lines). At log P_e = 15.68, the flux predicted by CHIANTI (v5.2) is a factor of 1.3 larger than that observed and a pressure of log P_e ≤ 15.30 would be required to fit the observed flux. Predicted flux ratios are in general more accurate than absolute fluxes, since they do not depend on abundances or ion fractions, so there may be small problems with the atomic data.

The corresponding lines of S XII are not observed.

4.6 Lines from iron ions

We now discuss the lines of iron according to their stage of ionization.

4.6.1 Fe IX to Fe XIV

For these ions, we rely on the Δn = 0 transitions observed with the EUVE, although the Fe X line is also observed with the LETGS. We also discuss the Fe XII forbidden lines that are observed with the STIS. At present, neither CHIANTI (v5.2) nor APED includes all transitions of the type Δn = 1 in these ions. We have updated earlier calculations of the EMLs by using CHIANTI (v5.2) (including those by Sim & Jordan 2003a, who used CHIANTI v4).

The lines of Fe IX to XIV used are all sensitive to n_e, either through the departure from Boltzmann populations in the levels of the ground term, or from the population of higher metastable states. Apart from the blended lines of Fe XIII at 203.83 Å the derived EMLs all increase with increasing n_e. The stronger 203.83-Å line ends on an excited level of the ground term and thus has the opposite behaviour. To give a smoothly increasing EMD from all the lines of Fe X to XV would require a value of log P_e ≤ 15.30; at higher pressures the Fe XIII locus lies below the mean value. Since there is no other evidence of such a low pressure, there might be small problems with the atomic data for Fe XIII in CHIANTI (v5.2) or in the EUVE line fluxes. With the inclusion of more levels in the atomic models of these ions, it might be possible to derive a value of P_e.

The resonance line of Fe XI at 171.07 Å leads to a relatively low EML, irrespective of pressures in the range from log P_e = 15.3 to 16.1. Even at log P_e = 16.10, the mean EMD leads to a flux that is larger than the value observed with the EUVE by a factor of 1.8. The 171-Å line falls near the long wavelength limit of the LETGS and the short wavelength limit of the EUVE medium-wavelength spectra; for both instruments there are significant uncertainties in the flux calibration at 171 Å and these could be one origin of the above discrepancy. As discussed by Laming et al. (1996), the ion balance calculations by Arnaud & Rothenflug (1985) lead to a larger EML than that found from the calculations by Arnaud & Raymond (1992). Near the peak emissivity, the difference between these calculations is a factor of 1.6, which would remove much of the discrepancy. However, the calculations by Arnaud & Raymond (1992) are expected to be more accurate. The final model can be used to estimate the line-centre opacity. This is close to 1; although scattering of photons out of the line of sight could occur, detailed radiative transfer calculations are needed to find the effect on the spatially integrated line fluxes.

The atomic data for Fe XIII have been revised since the work by Jordan et al. (2001a,b), who used CHIANTI (v3.01), and by Sim & Jordan (2003a), who used CHIANTI (v4.2). We have therefore re-examined the difference between the fluxes predicted by CHIANTI (v5.2) for the forbidden lines at 1242 and 1349 Å and the EUVE lines in the blend around 196 Å. For the lines at 1242 Å and at around 195 Å, Jordan et al. (2001a) found a difference of a factor of 3 between their EMLs. This is now reduced to a factor of 1.8 (at log P_e = 15.68) or 2.2 and 1.5 (at log P_e = 15.30 and 16.10, respectively). The small dependence on P_e arises from a small increase in the forbidden line fluxes, and a small decrease in the extreme-UV (EUV) line fluxes, with increasing P_e. Using the absolute line fluxes and the mean EMD, the agreement between the observed and predicted fluxes for the EUVE lines is very good (to within a factor of 1.1 over the above range of P_e) but the flux in the line at 1242 Å is predicted to be smaller than that observed, by the factors given above. Although differences in the fluxes arising from the different dates of the observations cannot be ruled out, neither can small corrections to the level populations for the forbidden lines (see below).

The ratio of the fluxes in the Fe XII forbidden lines at 1242 and 1349 Å is insensitive to log P_e over the range from about 15.0 to 16.0, but is useful in placing an upper limit on P_e. In ε Eri, the observed ratio is 1.88 (±0.2) (and other main-sequence stars show a similar ratio) (Jordan et al. 2001a). Using a single temperature of
log $T_e = 6.15$. CHIANTI (v3.01) leads to log $P_e = 15.72$, with an upper limit of 16.17. CHIANTI (v4.2) leads to log $P_e = 15.53$, with an upper limit of 16.07. However, CHIANTI (v5.2) leads to a pressure of log $P_e = 14.21$, which is much lower than the transition region pressure found in Section 3. The upper limit is 15.80, which is consistent with the transition region pressure (15.97 ± 0.20), but not with the pressure of log $P_e = 16.14$ found in the final model at log $T_e = 6.15$. At present, we suggest that the atomic data used in CHIANTI (v3.01 or v4.2) give a better fit to the forbidden line flux ratio than do those in CHIANTI (v5.2) (see Storey et al. 2005).

The value of the Fe xii forbidden line flux ratio provides a very sensitive test of the atomic data for these lines. For example, Jordan et al. (2001b) pointed out that the ratio of 2.7 predicted by Binello et al. (2001) could not be correct. It is also of interest to compare the population of the 3p32 P1 level from CHIANTI (v5.2) with that predicted empirically by Gabriel & Jordan (1975) on the basis of solar observations. At $T_e = 1.65 \times 10^6$ K and $n_e = 3 \times 10^8$ cm$^{-3}$, CHIANTI (v5.2) gives a level population (relative to that of the ion) of 2.9 $\times$ 10$^{-4}$, whereas the solar observations led to values between 3.3 and 5.1 $\times$ 10$^{-4}$. Thus, there is other observational support for a larger $2P_1$ population.

Given that the final EMD peaks around the temperature where Fe xv and xvi are formed, one might expect lines of Fe xiv to be present in the X-ray region. There are four weak lines around 76 Å that are also present in the LETGS spectra of Procyon (Raassen et al. 2002), α Cen A and B (Raassen et al. 2003) and Capella (Sanz-Forcada et al. 2004). In $\epsilon$ Eri, the lines are at 75.91, 76.04, 76.13 and 76.53 Å. We propose that the lines at 76.04 and 76.13 Å are due to the 3d $^2$D–4f $^2$F transitions in Fe xiv, but owing to the absence of these lines in CHIANTI or APED we cannot check this predicted empirically by Gabriel & Jordan (1975) on the basis of the population of the 3p32 P1 level from CHIANTI (v5.2) with that predicted by both CHIANTI (v5.2) and Keenan et al. (2006), which agree well with each other. This ratio is also lower than expected in Capella, but agrees with the theoretical value to within the uncertainties.

The 3s 4f $^1$F–3s 3d $^1$D transition occurs at 73.47 Å, but is potentially blended with a line of Ne viii at 73.48 Å. Interpreting the observed line flux as being due entirely to Ne viii, using CHIANTI and the mean EMD gives an observed-to-predicted flux ratio that is a factor of 4.0 too large. Assuming that the predicted Ne vii flux is correct, its contribution can be removed to give an Fe xv flux of 2.04 $\times$ 10$^{-14}$ erg cm$^{-2}$ s$^{-1}$ and a flux ratio of 0.68. The temperature sensitivity of the Fe xv 73.47-Å line is not given by Keenan et al. (2006) so that the temperature sensitivity of the triplet lines from the same configuration has been used to find the expected flux ratio at log $T_e = 6.50$. This corrected flux ratio agrees well with the flux ratio predicted by Keenan et al. (2006), but the flux ratio predicted using CHIANTI is larger. However, the stronger member of the Ne vii multiplet that should occur at 73.56 Å is not present with the flux expected from CHIANTI. For this reason, we have also found the observed flux ratio assuming no contribution from Ne vii, and this agrees with that predicted using CHIANTI, to within the uncertainties.

The predicted and observed flux ratios are summarized in Table 5. We have checked that the uncertainties in the observed ratios also cover the range of values derived using a range of background levels. Of the X-ray lines, only the observed flux ratio for the line at 69.68 Å in $\epsilon$ Eri is discordant with the calculations. On average, the flux ratios predicted by Keenan et al. (2006) are smaller than those predicted using CHIANTI (v5.2).

At present, the atomic model does not include levels with $n$ larger than 5; this could be one cause of the remaining differences between the observed and calculated flux ratios for transitions from the $n = 4$ levels. Also, from section 2.3 of Landi et al. (2006), the effects of recombination and ionization on the populations of excited states have not yet been included for Fe xv.

### Table 5. Fe xv line flux ratios, relative to the flux in the 59.40-Å line.$^a$

| Line (Å) | Keenan et al. (2006) | CHIANTI (v5.2)$^b$ | Observed |
|----------|---------------------|-------------------|----------|
| log $T_e$ | 52.91 | 6.3 | 0.65 ± 0.13 | 0.81 | 0.75 ± 0.17 |
| 6.5 | 0.64 ± 0.13 | 69.68 | 6.3 | 3.1 ± 0.6 | 2.75 | 1.85 ± 0.31 |
| 6.5 | 2.7 ± 0.5 | 73.47 | 6.5 | 0.97 ± 0.16 | 1.01 | 0.68 ± 0.20$^c$ |
| 6.5 | 0.75 ± 0.12 | 284.2 | 56.0 | 34.8 ± 5.6 |

$^a$This is 2.99 $\pm$ 0.44 $\times$ 10$^{-14}$ erg cm$^{-2}$ s$^{-1}$, when corrected for absorption in the ISM. $^b$Predicted using the mean EMD (see Table 10), using log $P_e = 16.10$. $^c$With the predicted contribution from Ne vii removed. $^d$Assuming no contribution from Ne vii.
4.6.3 Fe XVI

In addition to the 3p–3s transitions observed with the EUVE, lines from the 4p–3s, 4d–3p, 4s–3p and 4f–3d transitions are observed in the LETGS spectrum. Thus, it is possible to test the atomic data used in CHIANTI (v5.2); these were not updated from those used in CHIANTI (v4.2).

To make comparisons between the observations and the theoretical values from Cornille et al. (1997) and CHIANTI (v5.2), we use the 4s–3p line at 63.72 Å as the standard line. Using CHIANTI (v5.2) and the integrated fluxes, the predicted and observed ratios of this line and the lines at 335 and 361 Å agree to within 7 per cent.

The 4p–3s transitions at 50.35 and 50.55 Å are blended with the third order of the Fe XVII line at 16.78 Å and with the density-sensitive line of Si X at 50.52 Å. The plus-side spectrum was used to carry out the deblending, because the effective area varies rapidly in the minus-side spectrum. The fluxes in the third-order lines of Fe XVII at 17.05 and 17.10 Å have been measured and the ratio of the fluxes in these lines to that at 16.78 Å has been found from the first-order spectrum. Using integrated fluxes, it is found that the Fe XVII line at 50.35 Å contributes 71 per cent of the observed flux. At log $T_e = 6.5$, the theoretical ratio of the Fe XVII lines at 50.35 and 50.55 Å is 1.93, from Cornille et al. (1997), or 1.82, from CHIANTI (v5.2), using integrated fluxes. Hence, using CHIANTI (v5.2), Fe XVII contributes 51 per cent of the line at 50.54 Å. The discrepancy between the flux ratio for the 50.35-Å line is then a factor of 1.7, using Cornille et al. (1997), or 2.5, using CHIANTI (v5.2). These factors are significantly larger than the average found from other lines and these lines are not used in finding the mean EMD.

The 4d–3p lines at 54.72 + 54.76 and 54.14 Å are observed only in the plus-side spectra, but are relatively strong, clean lines. Using the calculations by Cornille et al. (1997), the flux ratio for the blended lines at $\approx 54.74$ Å is a factor of 1.3 smaller than that observed. Using CHIANTI (v5.2), at the same $T_e$, or using integrated fluxes, the predicted flux ratio is a factor of 1.9 smaller than that observed.

The 4f–3d transitions occur at 66.37 and 66.26 Å. The observed ratios of the fluxes in these lines to that of the 63.72 Å line are closer to the observed ratios than are those using CHIANTI (v5.2) and the mean EMD. Relative to the 4s–3p and 4f–3d transitions, the one line flux ratios do increase with $T_e$, but the mean EMD shows that little material exists at the very high temperatures required to bring the predicted ratios closer to those observed. Also, the lower fluxes measured for the higher ions of iron with the LETGS, compared with those measured from EUVE or the RGS, make it unlikely that flaring was present in the LETGS spectra.

As for Fe XV, the effects of including further $n$ states and of recombination and ionization to and from excited states still need to be investigated.

4.6.4 Fe XVII

The 2p$^3$3d–2p$^6$ transitions are observed at 15.01 and 15.26 Å and the 2p$^3$3s–2p$^6$ transitions are observed at 16.78 and 17.05 + 17.10 Å. The predicted flux in the 15.01-Å line is only slightly too large, in spite of the fact that the lower collisional excitation rate suggested by laboratory measurements by Brown et al. (2006) has not been adopted in CHIANTI (v5.2). The line at 15.26 Å is observed to be a factor of 1.45 stronger than predicted. Without knowing $n_e$ at high values of $T_e$, optical depth effects cannot be estimated. However, Brickhouse & Schmelz (2006) have suggested that an inner shell transition of Fe XVII occurs at 15.26 Å, but its contribution is not included in the predicted flux.

4.6.5 Fe XVIII

The resonance lines of Fe XVIII at 93.92 and 103.94 Å are observed with the EUVE and the LETGS, but the latter line is weak in both spectra. Since the ratio of its flux to that of the line at 93.92 Å does not agree with the theoretical value, the line at 103.94 Å is not used in determining the EMD.
Fe xvi has many transitions of the type Δn = 1 so there are few strong lines. The optimum temperature of line formation in a uniform plasma is \( \log T_e \approx 6.8 \), whereas the mean EMD peaks at around \( \log T_e = 6.6 \). Thus, the lines of Fe xvi are relatively weaker than in stars that have hotter coronae.

As discussed in Section 4.1, the O viii Lyman β line at 16.01 Å is blended with a transition of Fe xvi at 16.00 Å. The observed relative fluxes of the strongest lines at 14.21 + 14.26, 16.08 and 93.92 Å agree with those predicted by CHIANTI (v5.2) and the mean EMD to within a factor of 1.1.

4.6.6 Fe xix and xx

The observed lines of these ions lie between 100 and 130 Å where Sanz-Forcada et al. (2004) point out that there are suspected problems with the flux calibration. The noise levels are also large.

The line of Fe xix at 108.36 Å is weak, and that at 101.55 Å is barely above the noise level. The EML from the line at 108.4 Å shows that the mean EMD decreases rapidly at temperatures above about \( \log T_e = 6.7 \). The observed ratio of these two lines (which has only a very small dependence on \( n_e \) over the range of interest) is larger than predicted, probably because of difficulty in extracting a reliable flux for the weaker line.

The decrease in the EMD at higher temperatures is confirmed by the line of Fe xx at 132.8 Å. Lines from higher ions are not definitely observed. Lines that are formed above the mean coronal temperature (allowing for the extensions of their EMLs to lower temperatures) are likely to be formed in stellar active regions, by analogy with the behaviour of the solar EMD. Since we do not have values of \( n_e \) at these high temperatures, we cannot model the active region component.

5 RELATIVE ABUNDANCES AND THE MEAN EMD

5.1 Relative photospheric abundances

The photospheric element abundances derived for ε Eri by Abia et al. (1988), Drake & Smith (1993), Zhao et al. (2002) and Bodaghee et al. (2003) (and references concerning Fe therein), were discussed by Sim & Jordan (2005), in the context of relative abundances in the lower transition region. All find photospheric abundances of iron that are lower than the solar value. Abia et al. (1988) used early photospheric models, and the solar abundances they derive differ significantly from recent values; their results are not considered further here. The abundances derived by the other authors, relative to solar values, are given in Table 7. These include the more recent values from Allende Prieto et al. (2004) that were adopted by Wood & Linsky (2006). Both Bodaghee et al. (2003) and Zhao et al. (2002) used local thermodynamic equilibrium model atmospheres by Kurucz (1993). Bodaghee et al. (2003) used the solar abundances by Anders & Grevesse (1989) in their solar models, but Zhao et al. (2002) determined the differential abundances using their observations of the Moon, so their results have quite small uncertainties. Allende Prieto et al. (2004) also determined differential abundances using the same lines in the solar spectrum. The set of values by Zhao et al. (2002) is the most complete and we have adopted these as the initial values in deriving the EMLs, but also make comparisons with those by Allende Prieto et al. (2004), Sanz-Forcada et al. (2004) also made comparisons with the relative abundances by Zhao et al. (2002). To convert the differential photospheric abundances in ε Eri to absolute abundances, we have used the solar photospheric abundances recommended by Asplund, Grevesse & Sauval (2005). While the solar photospheric abundances adopted affect the stellar photospheric abundances, they do not influence the relative stellar coronal abundances discussed in Section 5.3.

The solar abundances recommended by Asplund et al. (2005) are based on exploratory three-dimensional modelling and are not directly comparable with the results of the one-dimensional stellar photospheric models. When comparing the stellar photospheric and coronal abundances in Section 5.3, we have also investigated results using solar abundances from Grevesse & Sauval (1998) that are also based on one-dimensional modelling. Photospheric abundances are not available for nitrogen or neon. For nitrogen, the mean of the carbon and oxygen differential abundances was used, whereas for neon, the solar abundance from Asplund et al. (2005) was initially adopted.

5.2 Derivation of the mean EMD

Fig. 3 shows the EMLs for lines from the H i-like, He i-like and Li i-like isoelectronic sequences (top panel), the EMLs from the lines of iron (middle panel) and all these lines (bottom panel). A value of \( \log P_e = 16.10 \) was adopted. These EMLs are apparent values, as defined in Section 6. Apart from the lines of Fe ix, Fe xvi and Fe xvi, the iron lines are formed over much smaller ranges of \( T_e \) than those from the above isoelectronic sequences. Because of these differences, the impression that the relative abundance of iron adopted is significantly too large is not correct. However, from the top panel it appears that the adopted relative abundance of nitrogen and silicon is quite accurate, that the relative abundance of carbon and magnesium is too large and that the relative abundance of nitrogen and oxygen is too small. From the bottom panel, it appears that the relative abundance of neon and iron is too small.

To adjust the relative abundances, the mean EMD must be found and used to recalculate the line fluxes; systematic differences in the calculated and observed fluxes for lines from the various elements can then be investigated. When attributing any differences to the effects of abundances, rather than to the shape of the EMD, it is assumed that the corrections to the abundances do not depend on \( T_e \).

The mean EMD to be derived is defined in terms of the value of the emission measure for a logarithmic temperature range of 0.30 dex, hereafter, EM(0.3). The value of \( P_e \) is assumed to be constant with \( T_e \), and the results using the three values \( \log P_e = 15.68, 15.30 \) and 16.10 have been investigated.

| [C/H] | [O/H] | [Mg/H] | [Si/H] | [S/H] | [Ca/H] | [Fe/H] | [Ni/H] |
|-------|-------|--------|--------|-------|--------|--------|--------|
| -0.10 (0.05) | -0.10 (0.07) | -0.10 (0.05) | -0.10 (0.07) | -0.09 (0.05) | -0.09 (0.05) | -0.16 (0.04) | Bodaghee et al. (2003) |
| -0.14 (0.05) | -0.12 (0.01) | -0.12 (0.01) | -0.12 (0.01) | -0.20 (0.03) | -0.20 (0.03) | Zhao et al. (2002) |
| -0.16 (0.02) | -0.16 (0.02) | -0.16 (0.02) | -0.16 (0.02) | -0.16 (0.02) | -0.16 (0.02) | -0.16 (0.02) | Drake & Smith (1993) |
| -0.01 (0.01) | -0.01 (0.01) | -0.01 (0.01) | -0.01 (0.01) | -0.01 (0.01) | -0.01 (0.01) | -0.01 (0.01) | Zhao et al. (2002) |
| -0.24 | -0.24 | -0.24 | -0.24 | -0.24 | -0.24 | -0.24 | Allende Prieto et al. (2004) |

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In order to find an initial EMD, the differences between the functions \( g(n_i, T_i) \) for the lines used must be taken into account; some lines are formed over a small range of \( T_i \), while others have contributions from a wide range of \( T_i \). The procedure developed by Jordan & Wilson (1971) and applied, with some modification, by Griffiths & Jordan (1998) is adopted. First, the fraction of the line formed over a temperature range of \( \log T_e = \log T_m \pm 0.15 \) is calculated, where \( T_m \) is the temperature at which the line emissivity has its maximum value, without allowing for any variation in the EMD with \( T_i \). The total contribution is found by integrating \( g(n_i, T_i) \) as a function of \( T_i \), cutting off the integration when the value is 0.01 times the maximum value of \( g(n_i, T_i) \). Thus, each locus of possible values of \( \text{EM}(T_i) \) is replaced by one value, referring to the range \( \Delta \log T_e = 0.3 \) about \( \log T_m \).

From the values of \( \text{EM}(0.3) \) derived from each line, it is then already clear that, relative to the lines of iron, the points for Ne \( \text{viii} \) lie above the mean, that is, the abundance of neon relative to iron is too small. At this point, the deblended fluxes of the lines of Ne \( \text{x} \) and Fe \( \text{xiv} \) discussed in Section 4.1 were used to derive a new starting abundance of 8.01 for neon. (See details in Section 5.3.) The individual values from each line have then been used to define the initial mean distribution of \( \text{EM}(0.3) \) with \( T_i \).

This initial mean distribution, defined at intervals of 0.1 in \( \log T_i \), is then iterated, taking into account the variations that we derive agree with those found by Wood & Linsky (2006) for the iron lines. Similarly, apart from those for oxygen and silicon, the abundances of the other elements were examined and small adjustments to the initial relative abundances adopted were made. We have found values relative to that of iron (as did Sanz-Forcada et al. 2004). The absolute values of the mean EMD will depend on the abundance of iron adopted. Sanz-Forcada et al. (2004) and Wood & Linsky (2006) derived absolute abundances of iron by measuring the line-to-continuum flux ratio, but we considered the continuum to be too weak for this to be useful; the error on the absolute abundance of iron found by Sanz-Forcada et al. (2004) is indeed quite large (7.20 \pm 0.2). Wood & Linsky (2006) find a value of 7.35, very close to the value of 7.33 that we adopt. The mean EMD derived using the final relative abundances and \( \log P_e = 16.10 \) is shown in Fig. 4 by the solid line with the grid points. Fig. 4 also includes the loci for lines of isoelectronic sequences omitted from Fig. 3 for the sake of clarity.

The mean EMD has also been derived using \( \log P_e = 15.30 \) and 15.68, adopting the above relative abundances. These agree with that derived using \( \log P_e = 16.10 \) to within mean values of \(-0.03 \) and \(-0.02 \) dex, respectively. The largest differences occur between \( \log T_e = 6.0 \) and 6.2, where they are \(-0.07 \) dex (for \( \log P_e = 15.30 \)) and \(-0.04 \) dex (for \( \log P_e = 15.68 \)). Thus, using a different value of \( P_e \) has an effect on the resulting mean EMD that is smaller than the uncertainties arising from the line flux measurements and the atomic data.

Alternatively, the relative abundances can be adjusted using the EMDs found with \( \log P_e = 15.30 \) and 15.68. For a fixed abundance of iron, these are on average 0.04 and 0.02 dex larger, respectively, than those derived using \( \log P_e = 16.10 \). Thus, the abundances derived (see Section 5.3) do not depend significantly on the value of \( P_e \) used in deriving the mean EMD.

5.3 Relative coronal abundances

In Table 8, we give the stellar photospheric abundances according to Zhao et al. (2002), scaled from the solar photospheric abundances of Asplund et al. (2005). The coronal abundances derived from the tables in Sanz-Forcada et al. (2004) and Wood & Linsky (2006) are also given. The values that we derive are given in the final column, on a scale where the coronal abundance of iron is set equal to the stellar photospheric value. We have investigated the likely uncertainties in the coronal abundances in several ways. Because we determine these relative to a fixed value of the iron abundance, using the full set of maximum observed fluxes, or the full set of minimum observed fluxes, has very little affect on the abundances derived. Instead, we give the uncertainties that result when the maximum fluxes for the lines of iron are used, together with the minimum fluxes of all other lines, and vice versa.

From Table 8, it can be seen that on the absolute scale adopted, all our abundances, except that for silicon, agree with those derived by Sanz-Forcada et al. (2004) to within our joint uncertainties. If their abundances are scaled to the photospheric abundance of iron (7.33), then all our abundances of elements used in finding the mean EMD agree with theirs, within the joint uncertainties. Similarly, apart from those for oxygen and silicon, the abundances that we derive agree with those found by Wood & Linsky (2006).

Although we do not include the possible lines of Ni \( \text{xi} \) in deriving the mean EMD, we give the mean abundance that results, since nickel is a low-FIP element. The lines of sulphur are all weak, and possibly blended, and individually give discordant abundances; the mean value is given. The possible line of Ca \( \text{xii} \) is predicted to be a factor of 4.2 weaker than that observed and, given the behaviour of the other low-FIP elements, there may be problems with the
identification, the calibration or the atomic data. The possible blend between the Ca xi and S xiv lines at 30.45 Å does not appear to support such a large abundance of calcium. Using the observed to predicted fluxes for the other two lines of S xiv suggests an abundance of calcium of 6.46. We do not include calcium in the discussions below.

Given that the starting abundances used by ourselves, Sanz-Forcada et al. (2004) and Wood & Linsky (2006) are different, and that we have used more up-to-date atomic data and independent flux measurements, these comparisons show that a line-based approach to the analysis of X-ray data gives reproducible results.

Figure 3. Emission measure loci for lines from the H I-like, He I-like and Li I-like isoelectronic sequences (top panel), lines of iron (middle panel), and all these lines (bottom panel). The photospheric abundances initially adopted for the transition region and corona of ϵ Eri are given in the bottom right-hand legend. A value of log $P_e = 16.10$ was adopted. The EMLs are apparent values – see Section 6.

Our EMD reproduces all the oxygen resonance lines well and the fit to the analysis of X-ray data gives reproducible results. Here, the final coronal abundances are used – see Table 8. A value of log $P_e = 16.10$ was adopted.

Table 8. Stellar photospheric abundances from Zhao et al. (2002), using Asplund et al. (2005) for the absolute scale; stellar coronal abundances derived by Sanz-Forcada et al. (2004), Wood & Linsky (2006) and from this work (the latter being scaled to the stellar photospheric iron abundance).

| Element | Zhao et al. (2002)$^a$ | Sanz-Forcada et al. (2004) | Wood & Linsky (2006)$^b$ | This work$^b$ |
|---------|-----------------------|---------------------------|---------------------------|--------------|
| C       | 8.33                  | 8.28 ± 0.18               | 8.15 ± 0.04               | 8.24 ± 0.09  |
| N       | (7.67)$^c$           | 7.74 ± 0.14               | 7.71 ± 0.04               | 7.82 ± 0.08  |
| O       | 8.50                  | 8.53 ± 0.04               | 8.46 ± 0.03               | 8.59 ± 0.04  |
| Ne      | (7.84)$^c$           | 8.09 ± 0.08               | 8.02 ± 0.04               | 8.06 ± 0.06  |
| Mg      | 7.39                  | 7.43 ± 0.10               | 7.46 ± 0.07               | 7.49 ± 0.13  |
| Si      | 7.35                  | 7.34 ± 0.07               | 7.33 ± 0.06               | 7.51 ± 0.07  |
| S$^d$   | 7.13                  | 7.21 ± 0.15               | 7.15 ± 0.26               | 7.29 ± 0.15  |
| Ca$^d$  | 6.20                  | 6.59 ± 0.20               | 6.70 ± 0.25               | (6.82 ± 0.14) |
| Fe      | 7.33                  | 7.20 ± 0.20               | 7.35                      | (7.33)       |
| Ni$^d$  | 6.13                  | 6.24 ± 0.11               | 5.98 ± 0.19               | 6.14 ± 0.08  |

$^a$The likely errors in the relative abundances are as given in Table 7. $^b$The errors are for the abundances relative to that of iron. $^c$Not given by Zhao et al. (2002); values initially adopted here – see the text. $^d$Not used in determining the mean EMD. $^e$The line involved is too weak to derive a reliable value.

However, the mean EMD that we derive differs from that found by Sanz-Forcada et al. (2004). Our EMD increases smoothly up to a peak value at around log $T_e = 6.6$, whereas that found by Sanz-Forcada et al. (2004) has two peaks, at log $T_e = 6.3$ and 6.75–6.8.

To test how well the EMD of Sanz-Forcada et al. (2004) reproduces the fluxes of the EUV lines of Fe x–Fe xii and Fe xiv included in our analyses, their volume EMD (that refers to intervals of 0.10 in log $T_e$) has been converted to the scale of our EMD over height (that refers to intervals of 0.30 in log $T_e$). The relative abundances derived are also examined for consistency between the various stages of ionization of a given element.

The EMD found by Sanz-Forcada et al. (2004) extends to only log $T_e = 5.7$ and so cannot account for the fluxes in the O vi lines. Our EMD reproduces all the oxygen resonance lines well and the
relative abundances found from these agree with each other to about a factor of 1.1. Similarly, apart from the resonance line of Fe x and Fe xiii lines, our mean EMD leads to the same relative abundance of iron, to within a factor of 1.3, for all lines used in ions up to and including Fe xviii. Sanz-Forcada et al. (2004) do not show observed to predicted fluxes for lines of ions between Fe xi and Fe xiv, so we assume that they did not include lines observed with the EUVE. We find that using their mean EMD gives relative iron abundances for these lines that depend on $T_e$ and span a range of a factor of 2.6. Also, the abundances that we find from Ne viii to Ne x are more self-consistent. Thus, overall, we consider that our mean EMD gives a better representation of the line fluxes and relative abundances. The reason why we derive similar mean relative abundances to those found by Sanz-Forcada et al. (2004) appears to be the influence of lines used in common in the higher temperature range where our mean EMDs are in closest agreement.

Wood & Linsky (2006) use a similar set of lines to those included by Sanz-Forcada et al. (2004), but adopt CHIANTI (v4.2) for the atomic data. The EMD that they find peaks at a similar temperature, but below $\log T_e = 6.0$ it is far smaller than ours, probably because they did not include the O vi lines observed with the FUSE.

Since there is interest in the possible presence of a FIP effect, in which elements with a low FIP (less than about 10 eV) have relatively larger abundances in the corona, we have examined the relative abundances of oxygen and iron in the corona and photosphere, since this ratio has the smallest uncertainty in the corona. We derive a coronal value of $\log(n_O/n_\text{Fe}) = 1.26 \pm 0.04$. Combining the photospheric abundances of Asplund et al. (2005) and the differential abundances by Zhao et al. (2002) or Allende Prieto et al. (2004) leads to stellar photospheric values of $\log(n_O/n_\text{Fe}) = 1.17$ or 1.23, respectively. Similarly, adopting the solar photospheric abundances of Grevesse & Sauval (1998) leads to values of 1.29 or 1.35. Thus, the largest difference between the photospheric and coronal relative abundances is $\pm 0.09$ in the logarithm, and a larger relative abundance of iron in the corona is found only when the solar photospheric abundances of Asplund et al. (2005) are adopted. On the basis of observations made with the EUVE, Laming et al. (1996) concluded that any FIP effect in $\zeta$ Eri was not significantly larger than that found in the solar corona. We conclude that there is no clear evidence for any FIP effect in the inner corona of $\zeta$ Eri.

There is also considerable interest in the relative abundance of neon to oxygen, given its relevance to models of the solar interior (Bahcall et al. 2005) and the difficulties caused by the lower neon to oxygen, given its relevance to models of the solar interior (cf. Asplund et al. 2005 and Grevesse & Sauval 2006). This ratio is significantly larger than the solar values of $-0.82$ and $-0.75$ recommended by Asplund et al. (2005) and Grevesse & Sauval (1998), respectively. It is also larger than the value of $-0.77$ found in the solar transition region by Young (2006), and is more similar to the mean stellar coronal value of $-0.39$ found by Drake & Testa (2005) (who used a more approximate method to find this ratio). The only obvious factors that could reduce the derived coronal abundance of neon would be the present limitations of the current atomic models for Ne viii and perhaps Ne x.

Abundances relative to that of iron can also be derived from deblended lines, but these do depend on the accuracy of the combined atomic data for the lines used. When the Ne x and Fe xvii lines at 12.14 Å are deblended, the derived flux in the Ne x line is a factor of 1.09 smaller than that predicted. If this difference is attributed to an incorrect neon abundance, then the abundance becomes 8.02, not 8.06. The predicted flux in the Fe xvii line at 12.29 Å is too small by a factor of 1.33, but we suspect that there are problems with the atomic data since the EMD based on the adopted abundance of 7.33 gives an overall fit to all the iron lines to within a smaller factor. Similarly, when the O viii and Fe xvii lines at 16.01 Å are deblended, the predicted flux in the O viii line is too small, but only by a factor of 1.16. The predicted flux in the Fe xviii line at 16.07 Å is also too small, but by only a factor of 1.05. Thus, to within the expected uncertainties, the relative abundances from these individual lines are consistent with those derived from the overall fits.

Table 9 gives the measured line fluxes, corrected for absorption by the ISM, and the observed-to-predicted flux ratios for the lines used in deriving the mean EMD (shown in bold face), based on our abundances given in Table 8. The flux ratios should be interpreted in light of Section 4, where blending and atomic data issues are discussed.

6 MODELS BASED ON THE ENERGY BALANCE

The method used has been set out by Jordan & Brown (1981) for plane-parallel geometry and for a spherically symmetric geometry by Pan & Jordan (1995). It has been extended by Sim & Jordan (2003a) to include emission from a restricted area at a given temperature and is only summarized here. Since the heating of the quiet corona occurs at heights much greater than the first pressure squared isothermal scaleheight over which the observed lines are predominantly formed, it is assumed that in the regions below, the divergence of the conductive flux is balanced by the local radiative losses. For a spherically symmetric atmosphere, with emitting area $A(r)$, and in hydrostatic equilibrium, the theoretical EMD is given by

$$\frac{d \log [\text{EM}(0.3)_{\text{lsf}}]}{d \log T_e} = \frac{2}{1} + \frac{2 \log P_e}{\log T_e} + \frac{d \log A(r)}{d \log T_e} \frac{2 P_{\text{lsf}}(T_e) \text{EM}(0.3)_{\text{lsf}}^2}{\kappa P_e T_e^2}.$$ (8)

Here, $P_{\text{lsf}}(T_e)$ is the radiative power-loss function and $\kappa$ is the constant in the coefficient of thermal conductivity (the small variations in $\kappa$ with $T_e$ are ignored here). Equation (8) can be solved iteratively in hydrostatic equilibrium, to provide the run of the gas and electron pressures, the radial height and EM(0.3)lsf with $T_e$.

We only observe the ‘apparent’ emission measure, given by

$$\text{EM}(0.3)_{\text{app}} = \text{EM}(0.3)_{\text{lsf}} G(r) f(r) \frac{A(r)}{A_e(r)},$$ (9)

where $f(r) = (r/R_e)^2$ and $G(r)$ is the fraction of the photons emitted that are not intercepted by the star,

$$G(r) = 0.5 \left(1 + \sqrt{1 - \left[\frac{1}{f(r)}\right]}\right).$$ (10)

When equation (8) is used, $A_e/A(r)$, the fractional area occupied at a given $r$, is set equal to 1.0.

In solving equation (8), it is first assumed that $G(r) = 1.0$ and $f(r) = 1.0$, and the apparent value of the EMD at a chosen peak coronal temperature are used as boundary conditions. The starting value of the total gas pressure $P_e$ is then found from these parameters and the isothermal pressure-squared scaleheight. The values of $G(r)$, $f(r)$ and the starting pressure are then updated in each iteration. Thus, the iterated solution also provides the calculated apparent emission measure $\text{EM}(0.3)_{\text{app}}$, which can be compared with that observed. The further boundary condition applied is that at the base temperature of...
log\(T_e = 5.3\), \(d \log \text{EM}(0.3)_{\text{app}}/d \log T_e = 0\), to fit the overall observed mean EMD, including the results from Sim & Jordan (2005).

If the solution is satisfactory, it will also reproduce the electron pressure of 15.97 ± 0.2 at log \(T_e = 5.3\).

Comparing EM(0.3)_{app} with EM(0.3)_{cal} allows any differences to be attributed to the effects of the emission originating mainly from a restricted area \(A(r)\) at a given \(T_e\). Additional terms can then be added to equation (8) to allow for \(A(r)/A(r)\) and its variation with \(T_e\). This results in equation (5) in Sim & Jordan (2003a), which gives the full expression for the gradient of log EM(0.3)_{cal} with respect to \(T_e\). The values of \(A(r)/A(r)\) can then be refined.

There are two differences between the application of the code used here and by Sim & Jordan (2003a). First, the maximum EMD is always associated with an isothermal region at the chosen coronal temperature; secondly, the form of \(P_{\text{rad}}(T_e)\) adopted is 2.80 \(\times 10^{-19}/T_e^{1/2}\), to take account of the additional lines now included in atomic data bases.

After investigating a number of models using equation (8), the one that gives the best fit to the constraints set out above has a coronal temperature of log \(T_e = 6.53\), a peak emission measure of log EM(0.3)_{cal} = 28.16 and a base pressure of log \(P_e = 15.97\). The EMDs derived from the line fluxes and this solution for EM(0.3)_{cal} are shown in Fig. 5 by the full and dotted lines, respectively. The area factors derived are given in column 4 of Table 10.

Because of the form of the energy balance equation and the boundary conditions chosen, the choice of the coronal temperature determines the coronal emission measure and pressure. The ratio of the base pressure (at log \(T_e = 5.3\)) to the coronal pressure has a constant value and the base pressure and temperature determine the theoretical value of the base EM(0.3), and hence the fractional area at the
base temperature. Thus, the parameter space of the models can be further explored without additional full calculations. The scaling laws that result will be discussed by Jordan & Ness (in preparation).

The apparent EMD is poorly determined between log $T_e = 5.3$ and 5.8, owing to the paucity of lines observed in this temperature range. When the derived areas are used in the full equation (5) of Sim & Jordan (2003a), the area is kept constant from log $T_e = 5.3$ to 5.8, at the value at log $T_e = 5.5$, where the lines of O VI are predominantly formed. This is justified by the small physical extent of this region. The resulting distribution of log [EM(0.3) cal $A_r(r)/A(r)$] as a function of log $T_e$ is shown in Fig. 5, and the new area factors that result are given in column 6 of Table 10. The radial extent and electron pressure in the final model are also given in Table 10. A further iteration with the new area factors was carried out to check that the new solution did not differ significantly from the previous one but, given the inherent uncertainties in the fluxes, atomic data and the constants used in the energy balance equation, a fully converged solution was not pursued.

The mean EMD derived directly from the observed line fluxes peaks at a temperature around log $T_e = 6.6 \pm 0.05$ and the lines of Fe xviii and to a lesser extent, Fe xvii, do appear to require material at higher temperatures. We attribute this emission to active regions, but cannot model them to remove their contribution, since the electron pressure is not known, and the energy balance used here will not be appropriate, since heating by other than thermal conduction is not included. Quiet coronal models with log $T_e \geq 6.6$ that satisfy the constraint on the EMD gradient at log $T_e = 5.3$ lead to base values of log $P_e$ that exceed the upper limit of 16.17 derived from the density-sensitive lines at about this temperature.

It is difficult to make detailed comparisons with other earlier work, but there are several early determinations of the temperature at which the peak EMD occurs. Giampapa et al. (1985) used observations of $\epsilon$ Eri made with the Imaging Proportional Counter (IPC) onboard the Einstein Observatory to deduce a single-temperature fit of log $T_e = 6.53$, close to the value of 6.60 ($\pm 0.05$) at which we and Wood & Linsky (2006) find the EMD to peak. However, the loop models that Giampapa et al. (1985) investigated had pressures that exceeded those found here and they remark that they could not simultaneously explain the IUE and X-ray spectra. Schmitt et al. (1990) also used these observations but could not find a single-temperature fit to the spectrum. The two-temperature fit that they suggested is not consistent with the present results. Because there was some suggestion that the absence of stars with coronal temperature ($T_e$) between log $T_e = 6.67$ and 6.88 in the sample studied by Schmitt et al. (1990), might arise from the energy response function of the IPC, Montesinos & Jordan (1993) used their scaling law between $T_e$, $g_e$ and the Rossby number, Ro, to predict the coronal temperature for $\epsilon$ Eri. With the currently adopted value of $g_e$, the scaling law predicts log $T_e = 6.57$, close to the value found here. The emission measures shown by Laming et al. (1996), based on observations with the EUVE, also peak at log $T_e = 6.5 \pm 0.1$. Thus, there is good agreement between the peak temperatures from the observations made with three different instruments.

The absolute scale of EM(0.3) app that we adopt depends on the value of the iron abundance used. If a smaller value were used, the values of EM(0.3) app would all increase. The intercept with EM(0.3) app would occur at a lower temperature and above this value, the area factors would exceed 1.0, which is not physically acceptable. Thus, it seems unlikely that the adopted value of the iron abundance is significantly too large. Conversely, using a larger iron abundance would lead to smaller values of EM(0.3) app. This would

Table 10. The apparent and calculated EMDs, the fractional areas derived, the radial extent above log $T_e = 5.3$ and $P_e$ in the final theoretical models.

| log $T_e$ (K) | log EM(0.3) app$^a$ | log EM(0.3) cal$^b$ | [AR]/$A_r$ $^c$ | log [EM(0.3) cal $A_r(r)/A(r)$]$^d$ | [AR]/$A_r$ $^e$ | (r–r0)/ (10$^5$ cm) | $P_e$ (10$^{16}$ cm$^{-3}$ K) |
|---------------|---------------------|---------------------|-----------------|--------------------------------|-----------------|-----------------|----------------|
| 5.30          | 26.700              | 27.281              | 0.26            | 27.453                         | 0.18            | 0.0             | 1.38           |
| 5.40          | 26.724              | 27.296              | 0.27            | 27.468                         | 0.18            | 5.4             | 1.40           |
| 5.50          | 26.725              | 27.343              | 0.24            | 27.515                         | 0.14            | 1.44 × 10$^4$   | 1.41           |
| 5.60          | 26.727              | 27.406              | 0.21            | 27.579                         | 0.14            | 3.04 × 10$^4$   | 1.43           |
| 5.70          | 26.736              | 27.479              | 0.18            | 27.653                         | 0.12            | 5.98 × 10$^4$   | 1.43           |
| 5.80          | 26.871              | 27.558              | 0.21            | 27.732                         | 0.14            | 1.15 × 10$^2$   | 1.44           |
| 5.90          | 27.061              | 27.641              | 0.26            | 27.816                         | 0.18            | 2.19 × 10$^2$   | 1.44           |
| 6.00          | 27.324              | 27.726              | 0.40            | 27.931                         | 0.25            | 4.29 × 10$^2$   | 1.43           |
| 6.10          | 27.586              | 27.812              | 0.59            | 28.136                         | 0.28            | 9.32 × 10$^2$   | 1.41           |
| 6.20          | 27.808              | 27.897              | 0.81            | 28.295                         | 0.33            | 2.16 × 10$^3$   | 1.37           |
| 6.30          | 27.928              | 27.981              | 0.89            | 28.402                         | 0.34            | 4.87 × 10$^3$   | 1.30           |
| 6.40          | 28.034              | 28.061              | 0.94            | 28.440                         | 0.39            | 1.03 × 10$^4$   | 1.20           |
| 6.50          | 28.121              | 28.138              | 0.96            | 28.487                         | 0.43            | 2.09 × 10$^4$   | 1.07           |
| 6.53          | 28.138              | 28.158              | 0.95            | 28.500                         | 0.43            | 2.58 × 10$^4$   | 1.02           |

$^a$Derived using log $P_e = 16.10$. $^b$Calculated with $A(r)/A_r(r) = 1.0$. $^c$Fractional areas calculated from columns 2 and 3. $^d$Calculated including a variable area factor. $^e$Fractional areas calculated from columns 2 and 5. $^f$Radial distance above $r = 5.18 \times 10^{10}$ cm at log $T_e = 5.30$. © 2008 The Authors. Journal compilation © 2008 RAS, MNRAS 385, 1691–1708
result in smaller values of $A(r)/A_\odot(r)$ by the same factor. Such solutions cannot be excluded.

7 DISCUSSION AND CONCLUSIONS

We have analysed line fluxes measured with the LETGS onboard Chandra to obtain an apparent EMD. As part of this work, we have examined the self-consistency of the results from lines of a given ion. There remain inconsistencies in the results from different lines of Fe XVI and, to a lesser extent, Fe XV. One source of these differences could be excitations to $n$ states not yet included in the atomic models, followed by cascades. Although the lines are weak, there is a significant difference between the observed and predicted fluxes for the $3p$–$2s$ transition in Ne VIII. The atomic models and data do not appear to have been updated in CHIANTI since v3. There also needs to be a proper treatment of recombination (including cascades) to the $1s 2s^3S$ level in the He I-like ions and di-electronic recombination needs to be included. The pressure indicated by the $f/i$ ratio in O VII is currently somewhat larger than expected from the final model. Although blends have been taken into account in analysing the Fe XIII lines at around 203.8 Å, the observed flux is lower than that predicted by CHIANTI (v5.2), unless $log P_e$ is lower than 15.30, which is not consistent with the results from other lines.

Line optical depths can be estimated using the final model that includes the variable area factors. Several lines, in particular the resonance line of Fe IX at 171 Å, are estimated to have line-centre optical depths approaching 1. The effects on the measured fluxes will depend on the geometry; scattering out of the line of sight would be expected for lines formed in supergranulation boundaries, but to find the expected flux when integrated over the whole star would require detailed radiative transfer calculations.

Relative element abundances have been determined in the upper transition region/corona. These agree with those found previously by Sanz-Forcada et al. (2004) to within the expected uncertainties, and quite well with those of Wood & Linsky (2006), in spite of differences in the mean EMDs. This reproducibility lends support to the individual line-based methods of deriving abundances. The EMD found here is based on line fluxes measured from the LETGS spectrum, the STIS spectrum (for O VII) and on the EUVE counts measured by Schmitt et al. (1996) for Fe IX to Fe XII and Fe XIV to Fe XVI. This EMD gives a consistent relative element abundance of iron for all stages of ionization included, apart from Fe IX, for which the line flux is observed to be weaker than predicted. Using the EMD found by Sanz-Forcada et al. (2004) leads to a larger difference between the abundances found from Fe XI, Fe XII and Fe XIV. On the basis of the relative abundances of oxygen and iron, we conclude that there is no clear evidence of any difference between photospheric and coronal abundances for low- and high-FIP elements. Indeed, none of our previous studies of stellar transition regions has shown any clear evidence of FIP effects. The spatially integrated X-ray line fluxes are dominated by the first pressure-squared isothermal scaleheight; if changes in relative element abundances are occurring at greater heights in the corona, they might not be detectable. (See also discussion by Wood & Linsky 2006 regarding correlations with stellar mass-loss rates.)

The important Ne/O relative abundance is found to be $-0.53$ (on a logarithmic scale), slightly smaller than the value of $-0.44$ found by Sanz-Forcada et al. (2004) and Wood & Linsky (2006), but larger than the values of $-0.82$ and $-0.75$ recommended by Asplund et al. (2005) and Greisser & Sauval (1998), respectively, for the solar atmosphere. It is also larger than that found for the solar transition region by Young (2006) ($-0.77$), Drake & Testa (2005) derived an even larger value ($-0.39$) from studies of a range of active stars, but used a more approximate method, not a full study of the EMD. Since the relative abundances of O and Fe found in the stellar corona agree with the range of stellar photospheric values to within $\pm 0.09$ in the logarithm, it appears that it is the abundance of Ne in the stellar corona that differs from the recommended solar values. Drake & Testa (2005) have suggested that a larger Ne/O abundance ratio could resolve the difficulties introduced by the adoption of the lower C, N and O abundances proposed by Asplund et al. (2005).

The similarity of the variation in the EMD with $T_e$ found by using lines from individual isoelectronic sequences or just one element can be seen from Figs 3 and 4. We are of the opinion that line-based analyses, using the full emission measure measure contribution functions, are the best method for determining the mean EMD.

Recent work by the other authors mentioned above has concentrated on deriving a mean EMD and element abundances but models of the atmosphere were not produced. We regard the main purpose of deriving the mean EMD is to use it in comparisons with theoretical models based on assumptions about the energy balance. Early work on modelling in terms of loop structures was not entirely successful (e.g. that by Giampapa et al. 1985). We intend to study a larger sample of stars in future work, to investigate the systematic behaviour with stellar activity.

The mean (apparent) EMD has been compared with the predictions of models based on the assumption of an energy balance between the divergence of the thermal conductive flux and the radiation losses. This allows the fractional area of the emitting material to be found as a function of $T_e$. When the variation in the fractional area is not included, we find a fractional area that is constant at about 0.24 up to $log T_e = 5.8$ and then increases with $T_e$ to reach about 1 by $log T_e = 6.4$. Allowing for the variation in the fractional area with $T_e$ increases the calculated values of $P_e$ and reduces the base fractional area to 0.14 and the coronal area factor to 0.43. The solutions derived here do not have the problem of fractional areas that are greater than 1 that occurred in the earlier analyses by Sim & Jordan (2005). The derived behaviour is similar to the trend in the area occupied by the supergranulation network boundaries in the solar transition region and inner corona, taking into account that the stellar transition region extends to higher temperatures, because the coronal temperature is higher. At the lower end of the temperature range, the surrounding material would be at near coronal temperatures; the coronal filling factor of less than 1 is consistent with the additional presence of active regions.

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