A *Chandra*/LETGS Survey of Main-sequence Stars

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Abstract

We analyze the X-ray spectra of 19 main-sequence stars observed by *Chandra* using its LETGS configuration. Emission measure (EM) distributions are computed based on emission line measurements, an analysis that also yields evaluations of coronal abundances. The use of newer atomic physics data results in significant changes compared to past published analyses. The stellar EM distributions correlate with surface X-ray flux ($F_X$) in a predictable way, regardless of spectral type. Thus, we provide EM distributions as a function of $F_X$, which can be used to estimate the EM distribution of any main-sequence star with a measured broadband X-ray luminosity. Comparisons are made with solar EM distributions, both full-disk distributions and spatially resolved ones from active regions (ARs), flares, and the quiet Sun. For moderately active stars, the slopes and magnitudes of the EM distributions are in excellent agreement with those of solar ARs for log $T < 6.6$, suggesting that such stars have surfaces completely filled with solar-like ARs. A stellar surface covered with solar X-class flares yields a reasonable approximation for the EM distributions of the most active stars. Unlike the EM distributions, coronal abundances are strongly dependent on spectral type, and we provide relations with surface temperature for both relative and absolute abundances. Finally, the coronal abundances of the exoplanet host star $\tau$ Boo A (F7 V) are anomalous, and we propose that this is due to the presence of the exoplanet.

Key words: stars: coronae – stars: late-type – X-rays: stars

Supporting material: machine-readable table

1. Introduction

Cool main-sequence stars are universally observed to be surrounded by hot coronae ($T = 10^6 - 7$ K), which represent the outermost atmospheric layers of such stars. The nature and origin of the surprisingly hot coronae has been a focal point of solar/stellar atmosphere research for decades. Stellar coronae are best studied at short wavelengths, as most of the emission from hot coronal material emerges as X-ray and extreme ultraviolet (EUV) radiation. Most of what we know about stellar coronae is based on soft X-ray observations from a series of spacecraft dating back to the 1970s (Güdel 2004; Güdel & Nazé 2009). This legacy continues today with the *Chandra* and *XMM-Newton* spacecraft, which have both been observing the X-ray sky since 1999.

Only so much can be learned about coronae from simple broadband flux measurements. A more detailed study requires high-resolution X-ray spectroscopy. Both *Chandra* and *XMM* carry gratings that provide spectra of unprecedented quality. The Reflection Grating Spectrometers (RGS) on *XMM* observe from 5 to 35 Å. *Chandra* carries two gratings, the High Energy Transmission Grating (HETG) and the Low Energy Transmission Grating (LETG). The former is generally paired with the ACIS-S detector to make the HETG Spectrometer (HETGS), which observes from 1.2 to 31 Å. The latter is generally paired with the HRC-S detector to make the LETG Spectrometer (LETGS), observing from 5 to 175 Å. With its particularly broad wavelength range, *Chandra*/LETGS is capable of observing significantly more lines of more atomic species than either *Chandra*/HETGS or *XMM*/RGS. Thus, we here focus on the analysis of *Chandra*/LETGS data.

Analysis of emission lines observed in a coronal X-ray spectrum involves the reconstruction of the coronal emission measure (EM) distribution. This process yields two crucial diagnostics. The first is the EM distribution itself, which describes the distribution of temperature in the corona, and the second involves the measurement of element abundances in the corona, which can be different from photospheric abundances. Both the temperature and abundance diagnostics are crucial for studying the mechanism(s) behind coronal heating.

As for the abundance diagnostic, the solar corona and wind exhibit an abundance pattern where elements with a low first ionization potential (FIP) have abundances that are enhanced relative to elements with high FIP (von Steiger et al. 1995; Feldman & Laming 2000; Laming 2015). Similar “FIP effects” have been observed for some stellar coronae (Drake et al. 1997; Laming & Drake 1999), but in other cases, coronal abundances appear to be close to photospheric (Drake et al. 1995; Audard et al. 2001; Sanz-Forcada et al. 2009). Finally, there are a number of cases where an “inverse-FIP effect” is observed, where low-FIP elements are depleted relative to high-FIP elements (Güdel et al. 2001; Huemoerder et al. 2001; Audard et al. 2003; Robrade & Schmitt 2005). On the Sun, an inverse-FIP effect has recently been detected for the first time near sunspots during flares (Doschek et al. 2015).

Significant effort has been expended to study how coronal abundances vary with activity level and spectral type. Initial attention focused on the connection between high activity and inverse FIP, as notoriously active M dwarfs and active binaries tend to have inverse FIP (Liefke et al. 2008). In a survey of early-G dwarfs, Telleschi et al. (2005) found an inverse-FIP or no FIP effect for the youngest and most active stars with ages less than 300 Myr, but solar-like FIP effects at older ages. The importance of spectral type for coronal abundances becomes clear if attention is focused only on main-sequence stars, particularly if extremely active stars with X-ray luminosities (in erg s$^{-1}$) of log $L_X > 29$ are ignored. For such stars, a surprisingly tight relation between FIP bias and spectral type.
is found, with M dwarfs having an inverse-FIP effect, which reduces toward no FIP effect at a mid-K spectral type, and then drifts toward a solar-like FIP effect for early G dwarfs (Wood & Linsky 2010; Wood et al. 2012). We refer to this relation as the “FIP-bias/spectral-type” (FBST) relation. This relation implies that all M dwarfs have inverse FIP, not just the very active ones. Thus, the vast majority of main-sequence stars in the Galaxy (all but the most active) are presumed to follow the FBST relation.

The element fractionation that occurs in the process of coronal heating is potentially a crucial diagnostic of this heating, and any successful coronal heating model should be able to explain the fractionation patterns that are observed on the Sun and other stars. Currently, the only theoretical framework for explaining both a FIP effect and an inverse-FIP effect involves the presence of ponderomotive forces induced by Alfvén waves passing through coronal loops, with the direction of the force along the loop depending on where the waves are introduced and where they reflect within the loops (Laming 2004, 2009, 2012; Wood & Laming 2013). If correct, this model would strongly support an important role for Alfvén waves (or their generation) in coronal heating.

Returning to the temperature diagnostic characterized by the EM distributions themselves, the coronal temperature is an even more direct diagnostic of coronal heating than the abundances, with higher temperatures implying more intense heating. Observations clearly show that more active stars, with higher X-ray luminosities, systematically have higher coronal temperatures, implying that increases in stellar activity are not simply a matter of filling the stellar surface with more and more identical active regions (ARs; Güdel et al. 1997; Schmitt 1997; Telleschi et al. 2005). The most recent empirical analysis is that of Johnstone & Güdel (2015), who find an impressively tight relation between mean coronal temperature and X-ray surface flux for cool main-sequence stars of all types: $T_{\text{cor}} = 0.11 F_X^{0.26}$, with $T_{\text{cor}}$ in MK units and $F_X$ in erg cm$^{-2}$ s$^{-1}$.

Such analyses rely on reducing the coronal temperature distribution to just one or two temperatures. In reality, however, coronal temperature distributions seem to be more or less continuous, as opposed to singly or doubly valued. With nearly 20 years of data now acquired by Chandra and XMM, it should be possible to use detailed EM distributions to more precisely describe how coronal temperature changes with increasing activity, thereby providing more detailed constraints for coronal heating models. For example, one interpretation of the increase in $T_{\text{cor}}$ with activity is that this is indicative of the increasing dominance of flare-like emission as stellar activity increases. Another interpretation is that it is indicative of the increasing prevalence of a population of hotter and presumably larger coronal loops (e.g., Güdel 2004; Reale 2014). A precise assessment of how the EM distribution evolves as activity increases could be helpful for distinguishing between such interpretations, and this is a central goal of our project.

To accomplish this goal, we conduct a survey of all main-sequence stars observed with Chandra/LETGS. We measure EM distributions using uniform analysis procedures, and assuming consistent atomic data. We also measure coronal abundances for our sample of stars, and study how they vary with activity and spectral type. This work overlaps strongly with work already done on the FBST relation (see above), but we expand this in a number of important ways. For example, the FBST relation defined above involves measurements of relative coronal abundances, particularly abundances of high-FIP elements relative to the best measured low-FIP element, Fe. However, we here also assess whether absolute abundances, for Fe in particular, vary with spectral type like the relative abundances.

A final motivation for this project is that our sample of Chandra/LETGS spectra includes two particularly noteworthy recent observations obtained by us. One is an observation of $\eta$ Lep (F1 V) in 2017 December, which is the earliest type main-sequence star successfully observed by Chandra with grating spectroscopy. This helps us to extend the FBST relation to earlier spectral types, and test whether stars with very thin convection zones have significantly different EM distributions. Moreover, in 2017 February–March, Chandra/LETGS observed the exoplanet host star $\tau$ Boo A (F7 V). This observation allows us to assess whether a very close-in, massive exoplanet can affect coronal temperatures and abundances in a way that makes the star clearly anomalous in our survey.

2. Sample Definition and Data Reduction

Our data sample is defined by existing Chandra/LETGS spectra of main-sequence stars deemed to be of sufficient quality for our purposes. Table 1 lists the 19 targets with spectra that have numerous enough detectable emission lines for an EM analysis to be performed. The targets are listed in order of spectral type, ranging from $\eta$ Lep (F1 V) to AD Leo (M4.5 V). The stars cover a wide range of activity levels, with X-ray luminosities (Column 6 in Table 1) ranging from $\log L_X = 26.99$ (in erg s$^{-1}$) for $\alpha$ Cen A to $\log L_X = 30.06$ for EK Dra and AB Dor A. These X-ray luminosities are for the canonical ROSAT PSPC soft X-ray bandpass of 0.1–2.4 keV (e.g., 5–120 Å), and are measured directly from the LETGS spectra themselves.

The fourth column of Table 1 lists radii for our stars, which are necessary to compute X-ray surface fluxes ($F_X$). Recent work provides evidence in favor of $F_X$ being a preferable measure of stellar activity compared with $L_X$ or $L_X/L_{bol}$ (Johnstone & Güdel 2015; Booth et al. 2017). Our default source of radius estimation is the relation of Barnes et al. (1978), but many radii are taken from more direct measurements (Kervella et al. 2003; Morin et al. 2008; Boyajian et al. 2012a, 2012b). The fifth column lists photospheric effective temperatures ($T_{\text{eff}}$). Most of these are from published spectral analyses of the individual stars (Valenti & Fischer 2005; Holmberg et al. 2009; Heiter et al. 2015). The remainder are estimated using the $T_{\text{eff}}$ versus $B − V$ relation of Valenti & Fischer (2005), or for stars later than early K, are taken from Table 5 in Pecaut & Mamajek (2013) or the $T_{\text{eff}}$ versus $V − K$ relation of Mann et al. (2015).

Our measured coronal abundances will have to be compared with photospheric abundances, so Table 1 lists photospheric abundance measurements for our stars. Following the usual convention, the abundances are listed logarithmically relative to solar photospheric abundances, so a value of 0.0 corresponds to an abundance equal to that of the solar photosphere. Throughout the paper, our default solar photospheric reference abundances are those of Asplund et al. (2009). For three companion stars (36 Oph B, $\xi$ Boo B, and 70 Oph B) we assume that the companion has the same abundances as the primary. No photospheric abundances are available for the two M dwarfs in the sample (AU Mic and AD Leo), as the formation of molecules makes photospheric abundance measurements very difficult. We
are forced to simply assume solar photospheric abundances for these stars. Likewise, photospheric abundance measurements are also very difficult for the two rapidly rotating stars EK Dra and AB Dor A because of line blending induced by rotational broadening. We once again simply assume solar abundances for these stars, consistent with crude estimates from optical spectra (Järvinen et al. 2007; Vilhu et al. 1987). Finally, for 61 Cyg AB, no measurements for C and N are available, so we simply assume values identical to Fe.

Table 2 lists the individual Chandra/LETGS observations that we have to work with, and we now describe the data reduction procedures we used to process the data. Rather than use the default processed spectrum, we process the data ourselves, using version 4.9 of the CIAO software provided by the Chandra X-ray Center (CXC; Fruscione et al. 2006). Many of our targets are binary stars with two separate resolved sources, and a tailored data processing is necessary in such cases anyway. An LETGS observation consists of a zeroth-order image of the target, with plus minus order X-ray spectra dispersed in opposite directions from the image. The zeroth-order image is used to establish the central reference point for a spectral extraction, and it is also useful for providing a broadband X-ray light curve for the observation, allowing flares to be identified. Although we do find some modest flares within the data, in all cases the flares are too weak or brief to contribute greatly to the counts of the overall integrated spectrum, therefore no attempt is made to remove any of the flares in the spectral extraction procedure.

Table 1
Main-sequence Stars Observed with Chandra/LETGS

| Star | Spectral Type | Dist. (pc) | Radius (R☉) | T eff (K) | log L_X (erg s⁻¹) | Relative to Solar |
|------|---------------|------------|--------------|-----------|-------------------|------------------|
| η Lep | F1 V | 14.9 | 1.56 | 6902 | 28.50 | –0.07 | –0.01 | 0.11 | –0.17 | –0.16 | –0.17 | –0.14 | 1 |
| π¹ Ori | F6 V | 8.07 | 1.32 | 6424 | 28.99 | 0.19 | 0.16 | –0.02 | 0.06 | 0.10 | 0.00 | –0.05 | 2 |
| τ Boo A | F7 V | 15.6 | 1.47 | 6387 | 28.76 | 0.33 | 0.30 | (0.33) | 0.33 | 0.33 | 0.33 | 0.22 | 3 |
| π¹ UMa | G1 V | 14.4 | 0.97 | 5768 | 28.99 | –0.11 | –0.04 | –0.26 | –0.09 | –0.11 | –0.25 | –0.24 | 2 |
| EK Dra | G1.5 V | 35.8 | 0.93 | 5845 | 30.06 | (0.00) | (0.00) | (0.00) | (0.00) | (0.00) | (0.00) | (0.00) | … |
| α Cen A | G2 V | 1.34 | 1.22 | 5792 | 26.99 | 0.18 | 0.25 | 0.39 | 0.32 | 0.40 | 0.12 | 0.20 | 2 |
| η Boo A | G8 V | 6.70 | 0.86 | 5570 | 28.91 | –0.10 | –0.09 | –0.26 | –0.10 | –0.35 | –0.26 | –0.24 | 2 |
| 70 Oph A | K0 V | 5.09 | 0.83 | 5202 | 28.09 | –0.10 | 0.03 | 0.09 | 0.18 | 0.09 | –0.05 | 0.06 | 2 |
| AB Dor A | K0 V | 15.2 | 0.79 | 5047 | 30.06 | (0.00) | (0.00) | (0.00) | (0.00) | (0.00) | (0.00) | (0.00) | … |
| α Cen B | K1 V | 1.34 | 0.86 | 5231 | 27.32 | 0.28 | 0.37 | 0.40 | 0.46 | 0.47 | 0.27 | 0.40 | 2 |
| 36 Oph A | K1 V | 5.99 | 0.69 | 5192 | 28.02 | –0.40 | –0.14 | –0.28 | –0.07 | –0.15 | –0.30 | –0.20 | 2 |
| 36 Oph B | K1 V | 5.99 | 0.59 | 5192 | 27.89 | –0.40 | –0.14 | –0.28 | –0.07 | –0.15 | –0.30 | –0.20 | 2 |
| ε Eri | K2 V | 3.22 | 0.74 | 5076 | 28.31 | –0.24 | –0.04 | –0.03 | –0.01 | –0.01 | –0.06 | –0.06 | 2 |
| η Boo B | K4 V | 6.70 | 0.61 | 4620 | 28.08 | –0.10 | –0.09 | –0.26 | –0.10 | –0.35 | –0.26 | –0.24 | 2 |
| 61 Cyg A | K5 V | 3.49 | 0.67 | 4374 | 27.03 | (–0.33) | (–0.33) | –0.18 | –0.29 | –0.36 | –0.33 | –0.39 | 4 |
| 70 Oph B | K5 V | 5.09 | 0.67 | 4450 | 27.97 | –0.10 | 0.03 | 0.09 | 0.18 | 0.09 | –0.05 | 0.06 | 2 |
| 61 Cyg B | K7 V | 3.49 | 0.60 | 4044 | 26.97 | (–0.38) | (–0.38) | –0.06 | –0.33 | –0.40 | –0.38 | –0.43 | 4 |
| AU Mic | M1 Ve | 9.91 | 0.61 | 3684 | 29.36 | (0.00) | (0.00) | (0.00) | (0.00) | (0.00) | (0.00) | (0.00) | … |
| AD Leo | M4.5 Ve | 4.89 | 0.38 | 3336 | 28.70 | (0.00) | (0.00) | (0.00) | (0.00) | (0.00) | (0.00) | (0.00) | … |

Table 2
Chandra/LETGS Observations

| Star | Obs. ID | Start Time | Exp. Time (ks) |
|------|---------|------------|---------------|
| η Lep | 20130 | 2017 Dec 15 21:12:58 | 116.3 |
| π¹ Ori | 20884 | 2017 Dec 11 17:03:23 | 38.0 |
| τ Boo AB | 12324 | 2010 Nov 9 12:17:00 | 57.2 |
| 70 Oph A | 13184 | 2010 Nov 21 19:23:36 | 19.9 |
| AB Dor AB | 17715 | 2017 Feb 27 22:41:08 | 47.9 |
| 61 Cyg AB | 20019 | 2017 Mar 4 23:43:24 | 28.5 |
| α Cen AB | 20020 | 2017 Mar 5 17:41:14 | 15.0 |
| η¹ UMa | 23 | 2000 Jan 15 6:14:22 | 30.0 |
| EK Dra | 1884 | 2001 Mar 19 3:20:38 | 65.5 |
| α Cen AB | 29 | 1999 Dec 24 10:38:20 | 79.6 |
| ζ Boo AB | 7432 | 2007 Jun 4 7:15:44 | 117.1 |
| 70 Oph A | 12332 | 2011 Jun 8 12:32:19 | 78.5 |
| AB Dor AB | 8899 | 2008 May 2 11:20:48 | 93.3 |
| 70 Oph A | 4482 | 2004 Jul 19 22:59:01 | 77.9 |
| AD Leo | 3762 | 2002 Dec 10 21:25:18 | 85.3 |
| 36 Oph A | 4483 | 2004 Jun 1 10:22:53 | 77.4 |
| ε Eri | 1869 | 2001 Mar 21 7:17:12 | 105.3 |
| 61 Cyg AB | 13651 | 2012 Feb 13 20:19:33 | 187.7 |
| AU Mic | 8894 | 2008 Jun 26 12:08:31 | 49.4 |
| AD Leo | 975 | 2000 Oct 24 15:06:16 | 48.1 |

Note.
* Values in parentheses assumed rather than measured.

References. (1) Yüce et al. (2011), (2) Allende Prieto et al. (2004), (3) Gonzalez & Laws (2007), (4) Jofré et al. (2015).

http://cxc.cfa.harvard.edu/ciao/threads/spectra_letghrs
spectra, we find it necessary to decrease Δs to minimize the background noise further. We use extraction windows as narrow as Δs = 12 pixels in some cases for this reason, although we always expand to Δs = 90 for wavelengths greater than 90 Å. The downside of narrow windows is a degradation of photometric accuracy, but for noisy spectra, this cost is more than balanced by the improvement in signal-to-noise ratio (S/N) due to the decreased background.

For binary stars it is necessary to avoid overlapping spectral extraction windows, so Δs can be limited by the stellar separation for close binaries. In such cases, it is generally possible to expand the extraction window in only one direction above 90 Å to avoid overlapping windows. There will be some degree of unresolved source blending at these higher wavelengths. We extract separate spectra for the two components of all binaries identified in Table 2 (e.g., the stars with “AB” in their names). In most cases, both components are considered in our target sample (see Table 1), with two exceptions, τ Boo B (M2 V) and AB Dor B (M5 V+M5-6 V). For τ Boo B and AB Dor B, we deem their spectra to be too noisy with too few detected lines for us to include these sources in our target list.

After background subtraction, the final step is to coadd the plus and minus orders. In many cases, it is necessary to shift either the plus or minus order spectra by up to four pixels before coaddition in order to line up the emission lines. This is an indication of the uncertainties in the LETGS wavelength calibration, which can vary in an unpredictable manner from observation to observation.

Multiple observations are listed in Table 2 for four stars. For η Lep, π Ori, and τ Boo, we coadd the observations to create a final spectrum. However, for α Cen AB the three observations are taken far apart in time, with the two stars at different points in their activity cycles. The α Cen system has been monitored regularly in X-rays by both Chandra and XMM, and using these data, Ayres (2014) estimates activity cycle periods of 19.2 ± 0.7 and 8.1 ± 0.2 year for α Cen A and B, respectively. For α Cen A, the first observation (ID 29) occurred near an activity cycle maximum in 1998, while the last two occurred closer to a minimum in 2008. We therefore consider two separate α Cen A spectra in our analysis, an α Cen A(hi) spectrum associated with observation ID 29, and an α Cen A (lo) spectrum that is a coaddition of the other two observations (IDs 7432 and 12332). Similarly, for α Cen B, the third observation is near an activity cycle maximum, with the other two near minima. Thus, we construct an α Cen B(hi) spectrum from observation ID 12332, and an α Cen B(lo) spectrum that is a coaddition of the other two observations (IDs 29 and 7432).

3. Line Identification and Measurement

The first step in our analysis of the Chandra/LETGS spectra is to identify and measure emission lines in the spectra. Our approach is to focus first on the highest quality spectra in our sample to establish the largest possible list of clearly detected and identified lines for the EM analysis, and then we search for only these lines in the other spectra. This is in effect a Bayesian approach to finding lines in the noisier spectra, as we only consider the precise wavelengths where the better quality data have informed us that we might reasonably find a line, which in turn allows us to be less conservative about claiming at least a marginal detection. A 1σ flux excess at exactly the right wavelength where a line is expected is far more likely to be a detection than a 1σ flux excess at some random location. Even for nondetections, we estimate upper limits for line fluxes, which are considered in the EM analysis described below.

The two high-quality spectra used in the initial line identification process are those of AB Dor A and α Cen B (lo). The former provides a representative spectrum of a high-activity star with high coronal temperatures, and the latter provides a representative spectrum of a low-activity star with low coronal temperatures. These spectra are shown in Figure 1.

We have previously published line lists and spectral analyses for a number of stars in our sample (Wood & Linsky 2006, 2010; Wood & Laming 2013), but we here consider far more lines, many of which are only identifiable now with better atomic data. Our primary line identification tool is version 7.1 of the CHIANTI atomic database (Dere et al. 1997; Landi et al. 2012, 2013). However, published line lists for the LETGS spectrum of Procyon are also valuable for identifying lower temperature lines (Raassen et al. 2002; Beiersdorfer et al. 2014). Table 3 provides our final list of 118 identified lines. These lines are also noted in Figure 1. A line formation temperature is estimated in the third column of the table, based on a mean temperature of the line contribution function. We also list in Table 3 counts measured for each line and for each star. Upper limits are measured for nondetections. These are 2σ limits computed from a conservatively broad 0.19 Å wavelength region around the line. Note that the print version of Table 3 is an abbreviated table listing only a few of the measurements for AB Dor A and α Cen B(lo). The full version of the table with all the line measurements for all of the stars is available online.

Given that we are only interested in lines that can be used in our EM analysis, Table 3 does not list any emission features that are blends of lines of different species. In order to be considered here, we have to believe that a line is at least ~80% from a single species, as ultimately verified using synthetic spectra computed from the EM distributions that we derive. Examples of blends visible in Figure 1 that are not considered are the O VIII+Fe XVIII blend at 16.0 Å and the Fe XX +Fe XXIII blend at 132.9 Å. We do naturally consider blends of lines of a single species, as we can use the line strengths in CHIANTI to divide the measured counts among the individual lines in the blend. Converting the line counts listed in Table 3 into line fluxes, which are also listed in the table (in units of photons cm$^{-2}$ s$^{-1}$), involves not only dividing the counts by exposure time and effective area, but also dividing the counts into the various individual lines in the case of blends. This is why, for example, a single count measurement is listed for the Si XIV 6.2 Å line, but two flux measurements are listed for it, as we have divided the counts between the λ rest = 6.180 Å and λ rest = 6.186 Å lines in the blend based on the CHIANTI emissivities.

Perhaps the most extreme single-species line blend is the Fe XX line at 12.8 Å, with dozens of Fe XX lines that are near that wavelength, which can be considered part of the blend. In such cases, our policy is to only list the two strongest lines in Table 3, and to only consider those two lines in the EM analysis. In the case of the Fe XX line, this means listing only the λ rest = 12.827 Å and λ rest = 12.845 Å lines, even though these two lines account for less than half the flux of the blend according to the CHIANTI emissivities. However, considering more than two lines in a blend would give the blend too much weight in the EM analysis, considering that a blended line is ultimately only a single detected emission feature. We
experimented with considering blends as only a single feature in the EM analysis, and found no significant change in our results. For the unresolved density-sensitive He-like triplets Si\text{XIII} \lambda 6.7 and Mg\text{XI} \lambda 9.2, we only list the strongest line in Table 3.

In identifying lines in the various spectra, it is necessary to be aware of higher order lines. In LETGS spectra, the higher orders are superposed on the first-order spectrum, although even orders are suppressed somewhat. For example, in the AB Dor A spectrum the features at 36.4 Å and 56.9 Å are third-order Ne\text{X} \lambda 12.1 and O\text{VIII} \lambda 19.0, respectively. The feature at 40.3 Å is not C\text{V} at \lambda_{\text{rest}} = 40.268 Å as Figure 1 might seem to suggest, but is instead third-order Ne\text{IX} \lambda 13.5, at least for AB Dor A.

A few lines are worthy of brief discussion, starting with the feature at 35.7 Å. In past analyses, using older versions of CHIANTI, we identified this as S\text{XIII} (Wood & Linsky 2006, 2010; Wood & Laming 2013). However, with CHIANTI version 7.1, this identification no longer seems to work. Nearby, there is a Ca\text{XI} line at 35.6 Å that is now in our line list, but we are no longer sure what the stronger 35.7 Å emission is, so it is not identified in Table 3 or Figure 1.

Another problematic line is at 94 Å. For the active stars that represent most of our sample, there is a line centered at 93.9 Å that is clearly Fe\text{XVIII} at \lambda_{\text{rest}} = 93.932 Å. However, for the inactive α Cen AB stars, there is instead a peak at 94.0 Å. This is clearly not Fe\text{XVIII}, although for α Cen B(hi) there is a marginal detection of Fe\text{XVIII} in the blue wing of the stronger 94.0 Å line. Given the numerous Fe\text{X} lines in the spectral region, the suspicion is that the 94.0 Å line is Fe\text{X}, but the version 7.1 CHIANTI line emissivities fail to provide an unambiguous identification, given that synthetic spectra fail badly to account for the line (see Figure 1). Testa et al. (2012) discuss the difficulties in modeling this feature in the LETGS spectrum of Procyon. This is an important feature for solar physics, as the Atmospheric Imaging Assembly (AIA) instrument on the Solar Dynamics Observatory (SDO) mission includes a filter bandpass centered at 94 Å, allowing the Sun to be monitored at this wavelength. Procedures for removing the low-temperature emission to determine a pure Fe\text{XVIII} image have been developed (Warren et al. 2012). These should work even though current line emissivities cannot reproduce disk-integrated spectra of this region from the Extreme ultraviolet.

Figure 1. Chandra/LETGS spectra of AB Dor A and α Cen B(lo), with the former representing a high-activity star with a high coronal temperature; and the latter representing a low-activity star with a low coronal temperature. Lines used in the EM analysis are identified in the figure. The red lines are synthetic spectra computed from the EM distributions derived from the spectra.
Variability Experiment (EVE) on SDO very well (Schonfeld et al. 2017).

Finally, one of the strongest lines in the $\alpha$ Cen B(lo) spectrum in Figure 1 is the Fe X line at 174.5 Å, which lies at the very end of the LETGS spectral range. However, the LETGS effective area is falling rapidly at this wavelength, and our experience suggests that the effective areas estimated at this wavelength by the CIAO data reduction software are unreliable. Thus, we cannot include this line in our analysis.

Figure 2 provides a graphical illustration of the various ionic species that are represented in our line list (e.g., Table 3), and how many of them are detected in our sample of stars. The detected lines cover a range of line formation temperatures from $\log T = 5.7$–7.4. For Mg and Si, every species from Mg VII–Mg XII and Si VIII–Si XIV is represented in the list. For the particularly important Fe sequence, every species from Fe VIII—Fe XXIV is represented, except for Fe XIII and Fe XXIII. Technically, there are even detected lines of Fe XIII and Fe XXIII at 76.5 Å and 132.9 Å, respectively (see Figure 1), but these are blends with lines of other species and so do not make our list. Needless to say, not every species is detected for every star. The species in red in Figure 2 are detected in nearly every spectrum, while those in green are only detected for a few.

A surprising amount can be learned from the line flux measurements simply by comparing the fluxes of different stars with each other, without the need for the complexity of the EM analysis that is described in the next section. With 21 spectra in our sample, many pairs of stars can be compared. Figure 3 shows examples of four such comparisons, plotting line flux ratios versus line formation temperature, with different colors indicating high-FIP ($\text{FIP} > 10 \text{ eV}$) and low-FIP ($\text{FIP} < 10 \text{ eV}$) elements. We exclude sulfur lines in these figures, as S is on the border between low FIP and high FIP. The flux ratios include the obvious corrections for differences in exposure time, distance, and radius; but there is no attempt to correct for different reference photospheric abundances. In cases where there is more than one line of a given species, we add fluxes of all lines detected for both stars before computing the flux ratio.

In all panels in Figure 3, the more active star with higher fluxes is divided by the less active star with lower line fluxes. Positive slopes are seen in each panel, demonstrating that coronal temperatures are higher for the more active stars. Separation between the low-FIP and high-FIP elements indicates a systematically different FIP effect. Higher line ratios are clearly seen for high-FIP elements in Figures 3(a)–(c). This could be said to indicate a weaker FIP effect for the more active stars, or...
Table 3
Line Measurements

| Ion   | λ_{line} (Å) | log T | AB Dor A Counts | AB Dor A Flux (10^{-5}) | α Cen B Counts | α Cen B Flux (10^{-5}) |
|-------|--------------|-------|-----------------|--------------------------|----------------|------------------------|
| Si XIV| 6.180        | 7.32  | 160.1 ± 36.7    | 2.95 ± 0.68              | <47.6         | <0.58                  |
|       | 6.186        |       |                 | 1.46 ± 0.33              | <46.2         | <0.54                  |
| Si XIII| 6.648       | 6.99  | 324.7 ± 42.6    | 5.40 ± 0.71              | 59.0 ± 24.5   | 29.47 ± 12.24          |
|       |              |       |                 |                          | 891.8 ± 43.7  | 366.38 ± 17.95         |

Note.

*Fluxes listed in units of 10^{-5} photons cm^{-2} s^{-1}.

This table is available in its entirety in machine-readable form.

Figure 2. Illustration of the ionic species with detected emission lines in our sample of main-sequence star Chandra/LETGS spectra. The species are plotted vs. line formation temperature. A color scale is used to indicate the number of stars for which a given species has a detected line.

Alternatively, a stronger inverse-FIP effect. The FIP bias seems to be relatively independent of temperature in the sense that we find no evidence of high-FIP ratios being higher than low-FIP ratios in one temperature range and lower in another temperature range. This is important because in the EM analysis described in the next section, we have to assume uniform abundances throughout the corona. Figure 3(d) compares the “(lo)" and “(hi)" spectra for α Cen B. The increase in coronal temperature with activity is very clear, but no evidence for any difference in FIP bias is apparent.

4. EM Analysis

With the line measurements made, the next step is an EM analysis to infer coronal temperature distributions and abundances. This analysis requires assumptions of collisional ionization equilibrium, Maxwellian velocity distributions, and uniform abundances throughout the corona. Two definitions of EM are used here, a volume EM, \( EM_V \) (in units of cm^{-3}), and a column EM, \( EM_C \) (in units of cm^{-2}). For the former,

\[
EM_V(T) = n_e^2 \frac{dV}{d \log T},
\]

where \( n_e \) is the coronal density and \( dV \) is a coronal volume element. This is the most natural expression of EM for unresolved, disk-integrated sources. Note that we express \( EM_V \) as a distribution in \( \log T \), rather than \( T \) as in some EM definitions. The observed flux for a given line can be expressed as

\[
f = \frac{1}{4\pi d^2} \int G(T, n_e)EM_V(T)d \log T,
\]

where \( d \) is the stellar distance and \( G(T, n_e) \) is the line contribution function, which includes both the line emissivity and the assumed elemental abundance of the atomic species in question.

For spatially resolved data (e.g., solar observations), a column EM is more intuitive,

\[
EM_C(T) = \frac{n_e^2 dh}{d \log T}.
\]

This EM version is also more useful when comparing EMs of stars with different radii, which we will be doing. The distance element \( dh \) can be related to \( dV \) by \( dV = 2\pi R^2 dh \), with \( R \) the radius of the star. We assume \( 2\pi R^2 \) instead of \( 4\pi R^2 \) because the emission from the back side of a star is hidden from us. Thus,

\[
EM_V = 2\pi R^2 EM_C.
\]

Knowing \( EM_V \) allows line fluxes to be computed with relative ease using Equation (2). However, the inverse problem of inferring \( EM_V \) from a set of emission line flux measurements is harder. For this purpose, we use version 2.97 of the PINTofALE software developed by Kashyap & Drake (2000), which includes a routine for computing EMs using a Markov chain Monte Carlo approach (Kashyap & Drake 1998). As in Section 3, version 7.1 of the CHIANTI database is the source of our line emissivities. The ionization equilibrium calculations of Mazzotta et al. (1998) are used, and we assume a typical coronal density of \( \log n_e = 10 \).

The PINTofALE routines include corrections for absorption from the interstellar medium (ISM). For stars as nearby as ours, ISM column densities are low and the corrections modest, but they are still well worth making for LETGS data, especially at higher wavelengths above 100 Å, where the effects of ISM absorption become more important. The logarithmic ISM hydrogen column densities assumed for our sample of stars, \( \log N_H \) (in units of cm^{-2}), are listed in Table 4. The primary source for these columns is a compilation of ISM column density measurements from Hubble Space Telescope (HST) spectra of the H I Lyα line (Wood et al. 2005). For π¹ UMa, the HST measurement is instead from Wood et al. (2014), while for
Figure 3. Line fluxes of various stars compared with each other by plotting line flux ratios vs. line formation temperature, with red (green) points indicating high-FIP (low-FIP) elements.

Table 4
Coronal Abundance Measurements: High-FIP Elements

| Star     | log N_H | [C/Fe] | [N/Fe] | [O/Fe] | [Ne/Fe] | [S/Fe] | [Ar/Fe] | F_{bias} |
|----------|---------|--------|--------|--------|---------|--------|---------|---------|
| η Lep    | 17.87   | 0.78±0.23 | ...    | 1.09±0.10 | 0.55±0.14 | ...    | ...    | −0.31±0.09 |
| π^3 Ori  | 17.93   | 0.70±0.10 | 0.11±0.18 | 0.92±0.05 | 0.46±0.05 | ...    | ...    | −0.41±0.07 |
| τ Boo A  | 17.68   | 0.83±0.26 | 0.69±0.35 | 1.15±0.15 | 0.50±0.34 | 0.11±0.38 | ...    | −0.21±0.09 |
| π^1 UMa  | 18.12   | 0.54±0.37 | 0.36±0.37 | 0.84±0.17 | 0.48±0.17 | 0.22±0.45 | ...    | −0.45±0.18 |
| EK Dra   | 18.08   | 0.95±0.21 | 0.19±0.63 | 1.12±0.11 | 0.96±0.07 | ...    | ...    | 0.00±0.13  |
| α Cen A(lo) | 17.71 | 0.69±0.09 | 0.06±0.18 | 0.35±0.15 | 0.41±0.09 | 0.07±0.57 | ...    | −0.54±0.30 |
| α Cen A(hi) | 17.61 | 0.61±0.10 | 0.25±0.12 | 0.67±0.11 | 0.33±0.13 | 0.12±0.12 | ...    | −0.50±0.15 |
| ζ Boo A  | 17.92   | 0.71±0.12 | 0.20±0.14 | 1.11±0.08 | 0.69±0.07 | 0.45±0.34 | 0.46±0.18 | −0.30±0.04 |
| 70 Oph A | 18.06   | 0.64±0.19 | 0.18±0.30 | 0.92±0.14 | 0.56±0.16 | ...    | ...    | −0.29±0.06 |
| AB Dor A | 18.29   | 1.50±0.09 | 1.06±0.08 | 1.75±0.03 | 1.35±0.05 | 0.14±0.06 | 0.18±0.07 | 0.60±0.08  |
| α Cen B(lo) | 17.61 | 0.70±0.04 | 0.14±0.14 | 0.87±0.05 | 0.33±0.13 | 0.52±0.27 | ...    | −0.38±0.15 |
| α Cen B(hi) | 17.61 | 0.70±0.05 | 0.08±0.10 | 0.89±0.03 | 0.44±0.08 | 0.76±0.23 | ...    | −0.36±0.09 |
| 36 Oph A | 17.85   | 0.57±0.47 | 0.44±0.47 | 1.19±0.13 | 0.88±0.11 | ...    | ...    | −0.14±0.10 |
| 36 Oph B | 17.85   | 0.44±0.37 | 0.33±0.64 | 1.10±0.19 | 0.67±0.24 | ...    | ...    | −0.27±0.10 |
| η Eri    | 17.88   | 0.90±0.01 | 0.42±0.06 | 1.20±0.02 | 0.84±0.03 | ...    | ...    | 0.06±0.07  |
| ζ Boo B  | 17.92   | 0.76±0.27 | ...    | 1.19±0.16 | 0.97±0.12 | ...    | ...    | −0.16±0.15 |
| 61 Cyg A | 18.13   | 0.95±0.24 | 0.34±0.40 | 1.12±0.19 | 0.79±0.39 | 0.20±0.43 | ...    | −0.01±0.04 |
| 70 Oph B | 18.06   | 0.96±0.41 | 0.70±0.33 | 1.35±0.18 | 1.01±0.14 | ...    | ...    | 0.15±0.10  |
| 61 Cyg B | 18.13   | 1.20±0.38 | ...    | 1.60±0.24 | 1.06±0.26 | ...    | ...    | 0.33±0.08  |
| AU Mic   | 18.36   | 1.63±0.09 | 1.19±0.09 | 1.77±0.06 | 1.39±0.05 | −0.20±0.09 | 0.68±0.13 |
| AD Leo   | 18.47   | 1.45±0.08 | 0.96±0.09 | 1.64±0.06 | 1.18±0.06 | 0.12±0.43 | ...    | 0.49±0.11  |
$\tau$ Boo A, log $N_H$ is estimated from the Mg II column density of Malamut et al. (2014) and the Mg depletion for this direction from Redfield & Linsky (2008). The EK Dra value of log $N_H = 18.08$ comes from a new HI Ly$\alpha$ measurement that we make from an archival HST spectrum (Ayres 2015). For details of how this type of analysis is done, see Wood et al. (2005). For $\eta$ Lep, $\pi^1$ Ori, and AB Dor A, no relevant ISM absorption measurements exist, so we assume the log $N_H$ values of stars close to these targets in the sky. Specifically, for $\eta$ Lep, $\pi^1$ Ori, and AB Dor A, the ISM columns listed in Table 4 are those measured toward HD 43162, $\chi^1$ Ori, and $\zeta$ Dor, respectively (Wood et al. 2005).

Figure 4 shows the EM$_V$ distributions derived by the EM analysis. The EMs are computed from log $T = 5.5$ to log $T = 7.8$, with a resolution of 0.1 dex. Error bars are 90% confidence intervals suggested by the Monte Carlo analysis. Technically, the line-based EM analysis only determines the shape of the EM distribution. A line-to-continuum ratio analysis is required to normalize it properly, which is described below. The temperature range in which the best-fit EMs are connected by a solid line provides an estimate of the temperature range that is constrained by detected lines, which is estimated from the peak temperatures of the line contribution functions, $G(T, n_l)$. The lower temperature bound is defined by the lowest peak temperature minus 0.3 dex, and the higher bound is the highest peak temperature plus 0.3 dex. Outside this range, the actual constraints on the EM are minimal. The best-fit EM values and the lower bounds indicated in the figure mean little outside this range, but the upper bounds are well constrained because the EM analysis does consider upper limits for all the undetected lines.

The EMs in Figure 4 are shown in order of increasing activity level, based on the logarithmic $F_X$ values indicated in the figure (in erg cm$^{-2}$ s$^{-1}$ units), which are computed from the radii and X-ray luminosities in Table 1. The general increase in both the magnitude of EM$_V$ and the mean coronal temperature with log $F_X$ is apparent. For $\alpha$ Cen A and B, both the “lo” and “hi” versions of the EM distribution are shown.

### 5. Coronal Abundance Measurements

In the EM analysis, the abundances of elements with lines of detected species are free parameters of the fits. However, the line-based analysis can only measure relative abundances, not the absolute abundances (i.e., the abundances relative to the dominant element, H). Given the prevalence of Fe lines in our spectra, Fe is the most obvious reference element to use for quoting relative abundances. Thus, the coronal abundances relative to Fe measured by the EM analysis are listed in Tables 4 and 5, with high-FIP elements listed in the former and low-FIP elements listed in the latter. We follow the common convention where abundances surrounded by square brackets are logarithmic, so the abundances in Tables 4 and 5 are listed in logarithmic form. Error bars are 90% confidence intervals, as in Figure 4.

The scattering processes that dominate continuum emission at X-ray wavelengths depend mostly on the abundances of H and He (e.g., Drake 1998), so the line-to-continuum ratio provides a measure of absolute abundances. With all coronal abundances measured relative to Fe from the emission line analysis, determining absolute abundances reduces to measuring the absolute abundance of Fe, i.e., [Fe/H]. We experimented with a number of different ways to perform the line-to-continuum analysis, using various wavelength regions. We even explored a sophisticated approach of allowing the considered wavelength regions to be different for different stars depending on where the highest continuum S/N is after line subtraction. However, this more complex approach did not seem to lead to any clear advantage, so for the sake of simplicity, we ultimately focus exclusively on the 25–40 Å region, which is relatively devoid of strong emission lines, except for C $\lambda\lambda$ 337.7, but is still at wavelengths short enough where the continuum is stronger.

Figure 5 shows the LETGS 25–40 Å spectra. Our procedure for measuring the continuum level here involves first measuring the total flux in this range (skipping a narrow wavelength region around the C $\lambda$ line), and then subtracting from this flux the integrated emission line flux inferred from a synthetic 25–40 Å spectrum computed using the EM distribution in Figure 4. This provides a continuum flux estimate, $F_{cont}$. Using the relevant PINToFALe procedures, we determine the absolute Fe abundance, [Fe/H], necessary to yield a continuum that reproduces $F_{cont}$. Figure 5 shows both the continuum-only and total line-plus-continuum synthetic spectra.

The inferred [Fe/H] values are indicated in Figure 5. The absolute Fe abundances are also listed in Table 5, but in the table we list them as [Fe/Fe$$_{\lambda}$], in order to indicate the coronal Fe abundance relative to the stellar photospheric Fe abundance from Table 1. For each $F_{cont}$ measurement, we also measure a Poissonian uncertainty based on the noise of the spectrum, and for eight of our spectra, we find $F_{cont}$ is less than 2σ above the noise. For these spectra, we conclude that we do not have a statistically significant detection of the continuum, meaning that we cannot make a meaningful measurement of [Fe/H]. For these eight cases, we have to assume [Fe/H] by other means. The assumed values are flagged in Table 5 and colored green in Figure 5.

For $\tau$ Boo A and $\pi^1$ UMa, the solution is to use published [Fe/H] measurements from XMM (Telleschi et al. 2005; Maggio et al. 2011). Observations from XMM are well suited for continuum measurement, with the high sensitivity of RGS and the possibility of considering spectra from both RGS and from the EPIC instrument. Pulse-height spectra from EPIC do not provide much spectral resolution, but this is far less important for continuum measurement than for line measurement. In contrast, LETGS is not ideal for continuum measurements, as the high background of the HRC-S detector can make it difficult to detect weak continua. For emission lines, this problem is mitigated somewhat by the excellent spectral resolution of LETGS, which means that the emission is isolated to a limited number of pixels with therefore a limited number of background counts. High spectral resolution is less helpful for detecting weak continuum emission, however.

This still leaves us with six spectra with no [Fe/H] measurement [61 Cyg A, 36 Oph B, 70 Oph a, $\eta$ Lep, $\alpha$ Cen A (hi), and $\alpha$ Cen A(lo)]. Estimates of some sort are required, as the [Fe/H] values are not only of interest in their own right, but are also necessary to normalize the EM distributions. Increasing [Fe/H] would correspond to lowering the EM$_V$ curves in Figure 4. We use our existing [Fe/H] measurements to estimate [Fe/H] for the stars without a measurement, but before describing this in detail, we first discuss the $F_{bias}$ values listed in the final column of Table 4.

The FIP-bias parameter, $F_{bias}$, represents an attempt to define a simple metric that can be used to quantify a star’s coronal
It is the average logarithmic abundance of four high-FIP elements (C, N, O, and Ne) relative to Fe, which is the best-constrained low-FIP element. This abundance ratio is computed relative to stellar photospheric abundances. Figure 6 shows explicitly how $F_{\text{bias}}$ is derived for our sample of LETGS-observed stars. In the figure, the abundance ratios of the six best-measured coronal abundances relative to Fe ($X/Fe$) from Tables 4 and 5 are plotted versus FIP (in eV), with a correction for the stellar photospheric abundance ratios, $[X/Fe]_p$, from Table 1. Thus, a value of 0.0 corresponds to a coronal abundance ratio identical to that of the stellar photosphere.

The abundance ratios in Figure 6 use the stellar photospheric abundance figures in Table 1, which must be converted into absolute abundances by assuming solar reference abundances. For that purpose, we use Asplund et al. (2009). No photospheric abundances are listed in Table 1 for N or Ne, as no stellar measurements exist for these elements. For N, we use O as a proxy, assuming $[N/Fe]_s = [O/Fe]_s$, with the N abundance ratios in Figure 6 reflecting this assumption. The situation for Ne is more complicated, as no direct photospheric measurements are possible even for the Sun, owing to a lack of photospheric Ne lines. Solar Ne abundances are instead inferred from coronal and transition region lines. The Ne abundance is often quoted relative to O, another high-FIP element. The Asplund et al. (2009) tabulation suggests Ne/O = 0.17, but much higher Ne/O values are generally found for stellar coronae. In particular, Drake & Testa (2005)
find Ne/O = 0.41. The most recent solar measurement from Young (2018) finds Ne/O = 0.24, which is a higher value than reported before, but still well below the stellar value. It remains debatable which measurement is preferable to represent the reference photospheric Ne abundance, but we follow past practice (e.g., Wood & Laming 2013) and use the Drake & Testa (2005) Ne/O ratio.

In Figure 6, \( F_{\text{bias}} \) is simply the average value of the four high-FIP elements, minus N in a few cases where no N lines are detected. This average is shown explicitly in the figure, with the stars displayed in order of increasing \( F_{\text{bias}} \). Values below 0.0 represent a solar-like FIP effect, while values above 0.0 represent an inverse-FIP effect. We compute a standard deviation of the four individual ratios, and use that as an estimate of the uncertainty in \( F_{\text{bias}} \). The \( F_{\text{bias}} \) values and their uncertainties are listed in Table 4.

At this point, we can compare our EM distributions and \( F_{\text{bias}} \) measurements with measurements made in the past, both from previous analyses of the same LETGS data studied here and from analyses of other data (e.g., XMM). We find two surprising differences. The first is that our new \( F_{\text{bias}} \) measurements are systematically higher than before. The mean and standard deviation of the difference is \( \Delta F_{\text{bias}} = +0.084 \pm 0.052 \). (See Table 2 of Wood et al. (2012) for a list of previous \( F_{\text{bias}} \) measurements.) The second difference concerns the shape of the EM distributions. A feature that we emphasized in the past was a surprisingly sharp peak in the EM distribution at log \( T \) = 6.6, which seemed ubiquitous and particularly convincing for high S/N data (Wood & Linsky 2010; Wood & Laming 2013). However, in Figure 4, such sharp and narrow peaks at log \( T \) = 6.6 do not exist. There is still generally an EM maximum near this temperature, but the exact log \( T \) = 6.6 temperature is no longer emphasized like it was before.

We trace both changes to large adjustments in the emissivities of Fe XVII lines in version 7.1 of CHIANTI, compared to older versions (Landi et al. 2012). There are four Fe XVII lines in the 15–17 Å range in our line list, which are usually the strongest Fe lines in the entire LETGS spectral range. These lines have rest wavelengths of \( \lambda_{\text{rest}} = 15.013 \) Å, 16.776 Å, 17.051 Å, and 17.096 Å; the last two being in an unresolved blend. The emissivity of the \( \lambda_{\text{rest}} = 15.013 \) Å line changed little from version 6 of CHIANTI, but the other three increased by an average of 60%. The strength of these Fe XVII lines makes it particularly important for an EM analysis to reproduce these line fluxes. With the line emissivities being apparently underestimated in the past, the EM reconstruction routine (PINToALE in our case) would correct for this in two ways: (1) increase the Fe abundance (thereby decreasing \( F_{\text{bias}} \)), and (2) create a sharp EM spike at log \( T \) = 6.6, which is the peak of the Fe XVII line contribution function. Further discussion of difficulties with Fe XVII emissivities can be found elsewhere (Beiersdorfer et al. 2004; Liang & Badnell 2010; Gillaspky et al. 2011; Bernitt et al. 2012). For our purposes, this serves as a reminder that uncertainties in line emissivities remain a crucial factor limiting the accuracy of EM analyses.

In Figure 7(a), \( F_{\text{bias}} \) is plotted versus spectral type, reproducing the FBST relation discussed in Section 1. In plots such as this, it is worthwhile to supplement our Chandra/LETGS measurements with other published results. Table 6 lists 10 stars with published X-ray spectral analyses, from which we can extract both \( F_{\text{bias}} \) and [Fe/Fe] values. These include the Sun, for which Schmelz et al. (2012) measure [Fe/Fe] = 0.33 and \( F_{\text{bias}} = -0.48 \). Table 6 also includes an LETGS \( F_{\text{bias}} \) measurement for GJ 338 AB. The GJ 338 AB spectrum was not of sufficient quality for a full EM analysis, but an \( F_{\text{bias}} \) value could be estimated from the Fe XVII/O VIII

### Table 5

#### Coronal Abundance Measurements: Low-FIP Elements

| Star   | [Mg/Fe] | [Si/Fe] | [Ca/Fe] | [Ni/Fe] | [Fe/Fe] |
|--------|---------|---------|---------|---------|---------|
| η Lep  | 0.16±0.11 | -0.02±0.11 | ... | -0.56±0.04 | (0.21) |
| δ Lep  | 0.23±0.10 | 0.24±0.07 | -1.12±0.27 | -0.09±0.06 | (0.07) |
| τ Boo A | 0.22±0.27 | 0.18±0.22 | ... | ... | (0.0) |
| η Lep  | 0.40±0.12 | 0.14±0.25 | ... | (0.24) | (0.12) |
| EK Dra  | 0.23±0.11 | 0.26±0.15 | ... | ... | 0.04±0.06 |
| α Cen A(lo) | 0.10±0.08 | 0.08±0.08 | -0.74±0.11 | -1.44±0.09 | (0.03) |
| α Cen A(hi) | 0.27±0.09 | 0.15±0.07 | -0.79±0.15 | -1.30±0.11 | (0.03) |
| ξ Boo A | 0.27±0.05 | 0.10±0.05 | ... | ... | 0.05±0.03 |
| 70 Oph A | 0.30±0.17 | 0.11±0.12 | ... | ... | (0.16) |
| AB Dor A | 0.18±0.13 | 0.11±0.05 | -1.06±0.03 | -0.64±0.01 | (0.01) |
| α Cen B(lo) | 0.26±0.06 | 0.03±0.10 | -0.91±0.18 | -1.15±0.11 | (0.01) |
| α Cen B(hi) | 0.26±0.07 | 0.11±0.04 | -1.02±0.04 | -1.16±0.07 | -0.61±0.05 |
| 36 Oph A | 0.63±0.18 | 0.19±0.21 | ... | ... | 0.25±0.15 |
| 36 Oph B | 0.19±0.41 | 0.01±0.35 | ... | ... | (0.16) |
| τ Cep A | 0.28±0.05 | 0.12±0.04 | -0.87±0.23 | -1.19±0.19 | (0.15) |
| ξ Boo B | 0.32±0.36 | 0.19±0.23 | -0.88±0.45 | -0.30±0.11 | (0.15) |
| 61 Cyg A | 0.16±0.44 | 0.15±0.36 | -0.99±0.49 | -0.14±0.03 | (0.34) |
| 70 Oph B | 0.55±0.25 | 0.24±0.23 | -0.61±0.25 | -0.31±0.16 | (0.25) |
| 61 Cyg B | 0.85±0.20 | 0.52±0.37 | ... | ... | 0.63±0.14 |
| AU Mic | 0.36±0.10 | 0.49±0.11 | ... | ... | 0.71±0.03 |
| AD Leo | 0.13±0.15 | 0.53±0.07 | -1.01±0.21 | -0.47±0.04 | (0.04) |

**Notes.**

a Assumed value based on the relation in Figure 7(c).

b Measurement from XMM/RGS data instead of Chandra/LETGS (Telleschi et al. 2005; Maggio et al. 2011).
Wood et al. 2012. The other measurements are either from XMM/RGS or Chandra/HETGS data. We expect all the published analyses (except for the Sun) will be affected by the aforementioned issue with the FeXVII line emissivities. Thus, based on our experience with the LETGS sample, we have adjusted the $F_{\text{bias}}$ values in Table 6 by $+0.084$.

As reported in past studies, we find that the relation between $F_{\text{bias}}$ and spectral type is surprisingly tight for main-sequence stars (see Section 1). However, this is the case only if one ignores particularly active stars with $\log L_X > 29$ (or $\log F_X > 7$). The three stars in our sample that fall in this category (EK Dra, AB Dor A, and AU Mic) are identified in red in Figure 7. In all three cases, the points lie well above the FBST relation. Evolved stars such as Procyon (F5 IV-V) and Capella (G8 III+G1 III) also seem to universally lie above the FBST relation (Wood & Laming 2013). The $\tau$ Boo A data point is also shown in red in Figure 7, as it seems slightly high.

Assessing whether the coronal abundances of $\tau$ Boo are affected by the presence of its close-in, massive exoplanet is one goal of our analysis (see Section 8).

In addition to the systematic changes in $F_{\text{bias}}$ noted above, there are two other notable changes to the FBST relation in Figure 7(a) compared to past work. One is the inclusion of error bar estimates for $F_{\text{bias}}$. The second and more important change is the extension of the relation to A and early-F spectral types. For the first time, we can definitively show that $F_{\text{bias}}$ values do not continue to decrease beyond early-G spectral types.

Figure 5. Highly smoothed LETGS spectra of the 25–40 Å region, which are used to measure the continuum level and determine absolute coronal abundances. The zero flux level is indicated by a horizontal black line. Blue lines indicate the inferred continuum level, based on the EM distributions in Figure 4. The red lines are the final synthetic spectra after the inclusion of the line emission. The absolute iron abundance suggested by the continuum level, [Fe/H], is indicated in each panel, with the panels arranged in order of increasing [Fe/H]. Values in green are assumed rather than measured, as these are cases where we do not have a statistically significant detection of the continuum.
The assessment of the FBST relation for earlier-type stars is possible thanks to the consideration of the very recent LETGS observation of $\eta$ Lep (F1 V) and the inclusion of the $XMM$ measurement for Altair (A7 V) from Robrade & Schmitt (2009). The primary reason Altair had not been considered before was that the spectral type of Altair is sometimes listed as A7 IV-V. Our experience with Procyon (F5 IV-V) suggests that luminosity class IV-V stars can be inconsistent with the FBST relation (Wood & Laming 2013). However, we have now concluded that we can consider Altair a main-sequence star for our purposes, primarily because the $F_{\text{bias}}$ measurements of Altair and $\eta$ Lep provide a consistent picture of an FBST relation that is flat at early spectral types (late-A through early-G). The surface gravity of Altair is significantly higher than that of Procyon, and closer to that expected for the main sequence (Malagnini & Morossi 1990). The mean radius for Altair of 1.93 $R_\odot$ might be considered consistent with an A7 V spectral type, considering the rapid rotation and resulting asphericity of this star (Ohishi et al. 2004). Peterson et al. (2006) conclude that Altair is close to the zero-age main sequence despite the IV-V classification.

Figure 7(b) shows another version of the FBST relation, with photospheric temperature replacing spectral type on the x-axis. This substitution makes it easier for us to quantify the FBST relation. We perform a linear fit to the $F_{\text{bias}}$ measurements (ignoring the red data points), with a flattening at a temperature that is another free parameter of the fit. The resulting relation is

$$F_{\text{bias}} = \begin{cases} 1.77 - 3.85 \times 10^{-4} T_{\text{eff}} & \text{for } T_{\text{eff}} < 5804 \\ -0.47 & \text{for } T_{\text{eff}} > 5804. \end{cases}$$

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The 1σ scatter of the \( F_{\text{bias}} \) values about the best fit is ±0.104. Figure 7(b) shows the 1σ and 2σ deviations from the best fit. The four red points that we have considered inconsistent with the FBST relation (albeit only tentatively for \( \tau \) Boo A) are indeed above the 2σ line, although a single green point (\( \eta \) Eri) is slightly above as well.

Switching from the relative abundances represented by \( F_{\text{bias}} \) to the absolute abundances represented by \( \text{[Fe/Fe]}_* \), Figure 7(c) plots \( \text{[Fe/Fe]}_* \) versus photospheric temperature. If the four red points are ignored, there seems to be a clear correlation, although with more scatter than the \( F_{\text{bias}} \) relations in Figures 7(a) and (b). A linear fit yields

\[
\text{[Fe/Fe]}_* = -1.29 + 2.18 \times 10^{-4} T_{\text{eff}}.
\]  

Unlike for the relative abundances, there is no evidence for a flattening of the relation at early spectral types. This conclusion, however, relies almost entirely on the Altair measurement, and suffers from our inability to detect the continuum and measure \( \text{[Fe/Fe]}_* \) for \( \eta \) Lep. We can now return to the issue of what to assume for \( \text{[Fe/Fe]}_* \) in the six cases where we do not have an X-ray continuum detection.

The individual element abundance measurements in Tables 4 and 5 can be perused to search for interesting behavior specific to particular elements. One example particularly worthy of note involves the coronal Si abundances of M dwarfs. In Figure 7(d), Si abundances relative to Fe are plotted versus \( T_{\text{eff}} \), with corrections for the reference photospheric abundances. For most stars, there is no dramatic difference between the coronal and photospheric Si/Fe ratio. The exceptions are the \( T_{\text{eff}} < 4000 \) K stars, i.e., the M dwarfs, which collectively have high coronal Si/Fe. The various panels of Figure 7 imply that for M dwarfs, low-FIP elements are
coronally depleted by a factor of 3–4, while the high-FIP element abundances are roughly photospheric. The exception is the low-FIP Si, which seems to behave more like a high-FIP element, with roughly photospheric coronal abundances.

One of the defining characteristics of M-dwarf photospheres is the formation of molecules, and for Si, the dominant molecule will be SiO (Tsuij 1973; Hauschildt et al. 1999). The other low-FIP elements that we are concerned with, Mg and Fe, are less inclined to form molecules, and the molecules that they do form (e.g., MgH and FeH) are more easily dissociated than SiO, which has a relatively high dissociation energy of 8.18 ± 0.17 eV (Reddy et al. 1998). We propose that the robustness of SiO leads to Si behaving more like a high-FIP element for M dwarfs, leading to the high Si/Fe ratios seen in Figure 7(d). Exploring this hypothesis further would require more detailed modeling than we can provide here.

Finally, the coronal Ne abundances are worthy of discussion because of their relevance for establishing the true cosmic abundance of Ne. Solar photospheric measurements are often used as reference “cosmic abundances” throughout astronomy, but there are no photospheric Ne lines, so the solar Ne abundance is instead estimated from transition region and coronal lines (Schmelz et al. 2005; Young 2005). Considering that Ne is one of the most abundant elements in the universe, establishing the proper cosmic abundance of Ne has broad ramifications. The Ne abundance is often measured relative to O, another high-FIP element. The reference solar photospheric abundances used here, from Asplund et al. (2009), assume Ne/O = 0.17 ± 0.05. However, stellar coronal measurements tend to be much higher than this, with Drake & Testa (2005) finding a weighted mean of Ne/O = 0.41 for a sample of 23 stars, with none of the stellar measurements as low as the solar one.

This raises two questions: (1) Why are the solar and stellar Ne/O measurements in disagreement, and (2) which value should be used to define the reference cosmic Ne abundance? As for the first question, Robrade et al. (2008) found evidence for lower stellar Ne/O ratios for less active stars, and argued that the low solar value may not be as inconsistent as the Drake & Testa (2005) sample suggests, since the sample is dominated by stars far more active than the Sun. In Figure 8, we look for an activity correlation for Ne/O within our sample of main-sequence stars. The Ne/O values in the figure are from Table 4, with uncertainties estimated from the standard deviation of the Ne/O values computed during the numerous Monte Carlo trials of PINTofALE. Excluding the anomalously high value for α Cen A(lo), we find a weighted mean of Ne/O = 0.39, in good agreement with Drake & Testa (2005). However, unlike Robrade et al. (2008), we see no evidence for any activity dependence, and like the Drake & Testa (2005) sample, not one of our measurements is as low as the solar one.

Nevertheless, the solar/stellar discrepancy has still improved somewhat, thanks mostly to the recent upward revision of the solar Ne/O ratio by Young (2018) to Ne/O = 0.24 ± 0.05 (see Figure 8). This measurement is from the quiet-Sun (QS) transition region. Landi & Testa (2015) find that Ne/O can be variable in the solar corona, with a high value of Ne/O = 0.25 ± 0.05 corresponding to the lowest activity corona, consistent with Young (2018). The most relevant comparison stars for the Sun in our sample are α Cen A and B. Our four measurements for those two stars are Ne/O = 1.15 ± 0.19, 0.46 ± 0.08, 0.29 ± 0.05, and 0.36 ± 0.04 for α Cen A(lo),
α Cen A(\text{hi}), \ α Cen B(\text{lo}), \text{ and } \ α Cen B(\text{hi}), \text{ respectively. The very high } \ α Cen A(\text{lo}) \text{ value is clearly anomalous, and we disregard it when computing the weighted mean quoted above. The } \ α Cen A(\text{lo}) \text{ spectrum is the only one in our sample where } Ne^\text{IX} \lambda 13.5 \text{ is undetected and } O^\text{VIII} \lambda 19.0 \text{ is just barely detected. The other } α Cen Ne/O \text{ values, however, are reasonably consistent with those of the more active stars in our sample. Note that Liefke & Schmitt (2006) measured } Ne/O = 0.28 \text{ from } XMM \text{ spectra of } α \text{ Cen AB combined, which will be dominated by the brighter } α \text{ Cen B. This is only slightly lower than our } α \text{ Cen B measurements. It remains an open question as to whether the cosmic Ne abundance is best assumed to be the new } Ne/O = 0.24 \pm 0.05 \text{ solar value from Young (2018), or our new stellar average of } Ne/O = 0.39. \text{ The solar value has the advantage of being from the transition region, where fractionation effects are believed to be less pronounced than in the corona. The stellar measurement has the advantage of coming from multiple sources.}

6. Disconcerting Absolute Abundance Measurements

The term “FIP effect” is often used to describe the coronal abundance anomalies seen in the solar corona. The $F_{\text{bias}}$ and $[\text{Fe}/\text{Fe}_*]$ quantities in Figure 7(b) and (c), however, represent two fundamentally different ways of thinking about the FIP effect, the former involving relative abundances and the latter involving absolute abundances. This can lead to significant confusion with regard to terminology. If it is said that a star has a solar-like FIP effect, does this mean low-FIP elements are enhanced relative to high-FIP elements, or does it mean that the absolute abundances of low-FIP elements are enhanced?

There is no distinction if high-FIP elements are not fractionated relative to H, which is after all a high-FIP element. The dashed line in Figure 7(c) shows explicitly what the $[\text{Fe}/\text{Fe}_*]$ curve should look like if high-FIP elements were rigidly tied to H, based on Equation (5). There is significant discrepancy from the observed relation. The comparison suggests that high-FIP coronal abundances are roughly photospheric for A and M stars, but are depleted for FGK stars. The Sun provides some evidence that high-FIP depletion is at least possible, with He in the slow solar wind being depleted by about a factor of two (von Steiger et al. 1995).

Figure 7(b) clearly shows that for solar-like G stars, low-FIP elements are enhanced relative to high-FIP elements, and that the stellar measurements are nicely consistent with those of the solar corona. In contrast, the linear fit in Figure 7(c) seems to imply that the absolute abundances of Fe in G star coronae are little different from photospheric, meaning that the relative abundance effect in Figure 7(b) is actually due primarily to depletion of high-FIP elements rather than an enhancement of low-FIP elements, in contrast to what is assumed to be the case for the Sun. In short, it is not clear that the G-dwarf absolute abundance measurements are consistent with the solar coronal measurements. The solar data point seems to be uncomfortably high in Figure 7(c), higher than the other G stars, and higher than all other stars except for Altair. The problem is even worse if the Schmelz et al. (2012) measurement of solar $[\text{Fe}/\text{Fe}_*]$ is replaced with the older but more canonical $[\text{Fe}/\text{Fe}_*] = +0.6$ value of Feldman & Laming (2000). This replacement would make the solar point in Figure 7(c) even more inconsistent with the stellar relation. So are we to say that early G stars have solar-like FIP effects, or not? In terms of relative abundances (e.g., Figure 7(b)), they definitely do, but in terms of absolute abundances (e.g., Figure 7(c)), they may not.

Both the significant scatter seen in Figure 7(c) and the questionable consistency between the solar and stellar measurements cast doubt on the accuracy of the absolute abundance measurements, which rely on the line-to-continuum ratio analysis. Possible pitfalls in this analysis have been discussed by Drake (1998) and Güdel (2004). A particularly important one involves the possibility that extensive blends of weak emission lines that are not in CHIANTI might be mistaken for the continuum. Is it possible that our continuum estimates in Figure 5 (blue lines) are too high as a result of a plethora of unknown, highly blended, weak emission lines in the 25–40 Å region? This would lead to underestimates of $[\text{Fe}/\text{Fe}_*]$, and if ubiquitous, could push the stellar values upward to be more consistent with the currently accepted solar value. Landi et al. (2013) provide an example of how numerous weak lines added in CHIANTI v7.1 significantly change and improve the appearance of synthetic spectra from 80 to 120 Å, when compared with CHIANTI v7.0. A similar change in the 25–40 Å region would have a dramatic effect on our line-to-continuum analysis.

It is worth noting that solar studies also have a history of ambiguity with regard to the issue of whether the solar coronal FIP effect is one of low-FIP enhancement or high-FIP depletion (Feldman & Widing 2003). Past claims of high-FIP depletions include Veck & Parkinson (1981), Fludra & Schmelz (1995), and Raymond et al. (1997). Current preference for the low-FIP enhancement interpretation comes in part from direct particle measurements of the slow solar wind, which are more suggestive of low-FIP enhancement and photospheric high-FIP, except for the aforementioned high-FIP He, with its factor of two depletion (von Steiger et al. 2000). Radio observations of the Sun also seem to provide support for an enhancement of roughly a factor of four of low-FIP elements in the corona (White et al. 2000; Schonfeld et al. 2015).

Direct comparison of the Sun with the most solar-like stars in our sample, α Cen A and B, provides further cause for unease with regard to the absolute abundance measurements. In Figure 4 the EM distributions of α Cen A and B are compared with solar distributions from Schonfeld et al. (2017) from full-disk SOHO/EVE spectra. The three solar distributions from Figure 9 of Schonfeld et al. (2017) represent a range of activity states for the solar corona. Significant variation in $EM_Y$ is observed only for $log T > 6.1$. This is very consistent with what we find when comparing the “(lo)” and “(hi)” EM distributions for α Cen A and B, where we see little variation for $log T < 6.1$.

Although α Cen A is generally considered the true Sun-like star of the α Cen binary because of its identical G2 V spectral type, α Cen B seems most similar to the Sun in terms of coronal properties. Not only are the $log F_X$ values measured for α Cen B the most solar-like, the shape of the α Cen B EM distribution is also the most similar to the Sun. In contrast, the α Cen A corona is cooler and has significantly lower $F_X$. Despite the coronal similarities, the magnitudes of $EM_Y$ are somewhat lower for the solar distributions than for α Cen B. This is mostly due to the different $[\text{Fe}/\text{Fe}_*]$ assumed in normalizing the $EM_Y$ curves. Schonfeld et al. (2017) simply assume $[\text{Fe}/\text{Fe}_*] = +0.6$ (Schonfeld et al. 2015), while for α Cen B we have measured identical $[\text{Fe}/\text{Fe}_*] = -0.61$ values from both the “(lo)” and “(hi)” spectra. The higher
reference photospheric Fe abundance for $\alpha$ Cen B ([Fe$_{\alpha}$/Fe$_{\alpha}$] = 0.27 from Table 1) moderates this discrepancy somewhat, but the coronal Fe abundance difference still adds up to 0.94 dex. The $\alpha$ Cen panel in Figure 4 shows the improved agreement between the intermediate solar EM$_V$ curve and the $\alpha$ Cen B(hi) distribution when the solar EM$_V$ curve is renormalized to assume the $\alpha$ Cen B coronal Fe abundance. Could it really be true that the Sun and $\alpha$ Cen B are coronally so similar in terms of X-ray flux and temperature distribution, but still exhibit radically different absolute abundance behavior, with the Sun having a substantial enhancement of Fe and $\alpha$ Cen B having a dramatic depletion?

Another coronal abundance comparison that can be made between the Sun and $\alpha$ Cen AB concerns the issue of activity cycle variability. Brooks et al. (2017) present evidence for a solar cycle variation of the FIP bias, with a stronger FIP effect at solar maximum. However, for both $\alpha$ Cen A and B, we find no significant difference in the $F_{\text{bias}}$ or [Fe/Fe$_{\alpha}$] measurements from the “(lo)” and “(hi)” spectra. This conclusion is evident even without an EM analysis, as shown in Figure 3(d). Thus, we conclude that $\alpha$ Cen AB coronal abundances do not vary significantly during the stellar activity cycles (Ayres 2014). The tightness of the FBST relation in Figures 7(a) and (b) by itself might imply little time variation of $F_{\text{bias}}$, considering that the stars will have been observed at different points in their various activity cycles. However, most of these stars are significantly more active than the Sun, and such stars tend to have more irregular activity cycles (Olsáh et al. 2016; Radick et al. 2018).

7. EM Variation with Activity

We now use the sample of EM distributions in Figure 4 to assess how the EMs of main-sequence stars vary with stellar activity. The Figure 4 distributions are provided in temperature bins with widths of 0.1 dex in $\log T$. For each $\log T$ bin, we can plot the measured EM values at that temperature versus $F_X$ for our sample of stars. Six of these plots are shown in Figure 9, for six different temperatures. We have converted from EM$_V$ into EM$_h$ to correct for the different radii of our stars. For each temperature bin, we find a reasonably smooth and consistent variation of EM$_h$ with $F_X$, with no evidence for any substantial spectral type variation. The M, K, and FG dwarfs seem to be consistent with each other. We fit third-order polynomials to the data points. Increases in EM$_h$ with $F_X$ are seen at all temperatures, but the relations are relatively flat for $\log T < 6.3$. The slope of the relation increases greatly for $\log T > 6.3$. It is remarkable that at $\log T = 6.3$, EM$_h$ for the extremely active EK Dra is not much higher than for the comparatively inactive $\alpha$ Cen B(hi), but at $\log T = 6.6$, the difference balloons to over three orders of magnitude.

The polynomial fits to the EM$_h$(T) versus $F_X$ relations can be used to define an average EM distribution as a function of $F_X$. This is done in Figure 10, which shows how a main-sequence
Figure 10. Main-sequence star EM distributions as a function of $F_X$, based on relations such as those in Figure 9. These are compared with an average solar active region distribution from Warren et al. (2012), a QS distribution from Kamio & Mariska (2012), and an average solar flare distribution from Warren (2014). The first two are from spatially resolved data, so these are true line-of-sight column emission measures, while the flare distribution is from disk-integrated data, so the EM$_h$ values are derived assuming the emission is uniformly spread over the visible surface.

The stellar EM$_h$ distribution varies with $F_X$ for log $F_X$ = 3.9–7.5. These curves are also provided in Table 7. This information can be used to estimate an EM distribution for any main-sequence star that only has a broadband X-ray flux measurement. It can also be used as the basis for future theoretical coronal heating models seeking to describe how heating changes with increasing stellar activity. If mean coronal temperatures are computed from these curves as prescribed by Johnstone & Güdel (2015), we can reproduce their power-law relation between $T_{\text{cor}}$ and $F_X$. Specifically, with $T_{\text{cor}}$ in MK units, we find $T_{\text{cor}} = 0.16F_X^{-0.24}$, compared with Johnstone & Güdel’s $T_{\text{cor}} = 0.11F_X^{0.26}$.

The stellar EM distributions in Figure 10 are compared with three solar distributions, representing the QS, solar active region (AR), and flaring (FL) Sun. The QS distribution is based on spectra from the EUV Imaging Spectrometer (EIS) on Hinode (Kamio & Mariska 2012). The AR distribution is an average of 15 ARs studied by Warren et al. (2012) using EIS and SDO/AIA data. Finally, the FL distribution is an average distribution from the peaks of 21 strong flares studied by Warren (2014) using SDO/EVE spectra.

In solar physics, decades of flare monitoring with the Geostationary Operational Environmental Satellite (GOES) spacecraft have led to the widespread use of the GOES system’s classification of flare strength, with the A, B, C, M, and X classes representing increasing decades of flare luminosity in the 1–8 Å bandpass. The 21 flares used to define the FL distribution range from M9.3 to X6.9. The individual flare EMs are shown explicitly in Figure 11(a), where we focus only on the two-minute interval of maximum integrated EM. In addition to using them to compute the average FL EM for Figure 10, we also compute synthetic spectra from them, and from those spectra, we estimate X-ray luminosities in the 5–120 Å bandpass most familiar to stellar astronomers. In Figure 11(b), these X-ray luminosities are plotted versus the GOES X-flare classification, $X_f$ (e.g., for an X5.2 flare, $X_f = 5.2$; for an M7.5 flare, $X_f = 0.75$). Not surprisingly, log $L_X$ increases with flare strength, with the fit to the data suggesting log $L_X = 0.56\log X_f + 27.75$. A typical X-flare has log $L_X \approx 28.0$, which is about five times the Sun’s quiescent X-ray luminosity, but is less than the quiescent emission from most of our stars (see Table 1).

There is an important distinction between the QS and AR distributions and the FL distribution in Figure 10. The former are from spatially resolved observations, so the EM$_h$ values are line-of-sight column EMs. Thus, these curves are indicative of the actual intensities of QS and AR regions. In contrast, the FL measurements are from disk-integrated SDO/EVE spectra, so EM$_h$ is computed in the same manner as for the stars, namely by computing EM$_h$ and then converting into a column EM$_h$ by dividing by the visible surface area of the star. In order to scale the FL EM$_h$ curve to indicate the actual flare intensity, it is necessary to divide the curve by the fractional surface coverage of a flare. For the 21 flares from Warren (2014), we estimate an average surface coverage of 0.3%. This corresponds to increasing the FL curve in Figure 10 by 2.5 dex.

We can compute synthetic spectra and X-ray luminosities for each of the solar EM distributions in Figure 10, and for the Schonfeld et al. (2017) full-disk solar EM distributions in Figure 4. We find log $L_X = 26.40, 28.66, and 27.94$ for the QS, AR, and FL distributions, respectively, while the Schonfeld et al. (2017) distributions imply a range of log $L_X = 26.86–27.40$. The QS value represents an estimate of the minimum possible X-ray luminosity for a solar-like G star. To generalize to main-sequence stars with different radii, this corresponds to a minimum surface flux of log $F_X = 3.61$. This is only about a factor of two lower than the lowest activity spectrum in our sample, that of α Cen A(lo). We are not aware of any main-sequence stars with detected X-ray fluxes below this limit (e.g., Schmitt & Liefke 2004). In principle, lower X-ray luminosities might be expected for very low metallicity stars, due to weak emission lines. However, the models of Suzuki (2018) suggest that such stars may end up brighter in X-rays because of much higher coronal densities.

Judge et al. (2017) recently found a very low upper limit for the X-ray luminosity of 16 Cyg B (G3 V), log $L_X < 25.5$, which is much lower than the minimum QS level of log $L_X = 26.4$ just quoted in the previous paragraph. However, the 16 Cyg observations are Chandra ACIS-I data, which cover an energy range of 0.3–2.5 keV, compared to the more traditional 0.1–2.4 keV ROSAT/PSPC range that we are using. These ranges seem similar, but the similarity is illusory for very inactive stars, as an inactive star will emit far more flux in the 0.1–0.3 keV bandpass than it will at >0.3 keV because of the very low coronal temperatures. At >0.3 keV, the emission of an extremely inactive star will be mostly from the O VII triplet at 22 Å and the C VI line at 33.7 Å. For an energy range of 0.3–2.5 keV, we find log $L_X = 25.37$ and log $F_X = 2.59$ for the QS distribution, a full order of magnitude lower than the 0.1–2.4 keV values quoted above. This 0.3–2.5 keV QS luminosity is not inconsistent with the very low 16 Cyg B limit of Judge et al. (2017). The luminosities inferred from the Schonfeld et al. (2017) distributions decrease to log $L_X = 26.15–26.96$ in the 0.3–2.5 keV range, which is more consistent with the α Cen AB luminosities quoted by Ayres (2014), which are for 0.2–2 keV.

It is impressive that the stellar EM$_h$ distributions in Figure 10 match the solar AR distribution so well for log $T < 6.6$ in terms
of both magnitude and slope of the distribution. This is true even though the solar and stellar analyses rely on completely different sets of emission lines measured from completely different wavelength regions. Between \( \log F_X = 5.0 \) and \( \log F_X = 6.5 \), the EM \( \beta \) curves basically lie right on top of each other for \( \log T < 6.4 \), which is suggestive of a saturation effect. This apparent saturation at close to the EM \( \beta \) level of the solar AR distribution strongly supports the idea first suggested by Drake et al. (2000) that the EM distributions of intermediate-activity stars like \( \epsilon \) Eri and \( \xi \) Boo A are nicely explained by the stellar surfaces being completely filled with solar-like ARs. The EM \( \beta \) values are very linear for \( \log T < 6.6 \), so we can assume \( \dot{E}_M(T) \propto T^{\beta} \) and measure the slope, \( \beta \). For the stellar EMs with \( \log F_X = 5.5-6.5 \), we find a mean slope of \( \beta = 2.50 \), in excellent agreement with the solar AR slope of \( \beta = 2.41 \). The EM seems to start increasing again for \( \log F_X > 7.0 \), but there are only a few stars in this high-activity regime, so the reality of this increase is questionable.

For the solar AR distribution, EM \( \beta \) drops precipitously for \( \log T > 6.6 \). It is therefore clear that while solar-like ARs may explain the stellar EM \( \beta \) for \( \log T < 6.6 \), they cannot account for the high-temperature EM \( \beta \) at \( \log T > 6.6 \), which becomes quite strong for stars with \( \log F_X > 5.5 \). For this emission, the only solar analog would be solar-like ARs. The EM \( \beta \) are nicely sensitive line ratios generally find electron densities (in units of \( \text{cm}^{-3} \)) of \( n_e \) \( \sim 10 \) for \( \log T < 6.6 \), consistent with typical solar AR densities. However, much higher densities of \( n_e \) are generally observed at \( \log T \sim 7 \) (Maggio et al. 2004; Testa et al. 2004; Osten et al. 2006; Liefke et al. 2008).

It is worth looking for abundance signatures that might also indicate a distinction between the \( \log T \sim 6.5 \) and \( \log T \sim 7.0 \) plasma. Unlike the quiescent solar corona, solar flares generally do not exhibit a FIP effect. This is demonstrably the case for the 21 SDO/EVE solar flares we have used to define the FL distribution in Figure 10, which have roughly photospheric abundances (Warren 2014). However, it is very hard for us to look for a FIP effect specifically at \( \log T \sim 7.0 \), because the only high-FIP line we have that is at least nominally formed at \( \log T > 6.6 \) is Ne X. This is illustrated best by Figure 3, which does not suggest dramatic differences in FIP bias at Ne X temperatures compared to lower temperatures. However, even Ne X is problematic, because the contribution function for this species is quite broad in temperature, meaning that if EM \( \beta \) is higher at \( \log T = 6.5 \) than at the nominal peak at \( \log T \approx 6.9 \) by as little as a factor of 3 or 4, then much of the Ne X emission will actually be coming from \( \log T \approx 6.5 \). Most of our stellar EM distributions do in fact peak near \( \log T = 6.5 \).

### Table 7

| \( \log T \) | 4.0 | 4.5 | 5.0 | 5.5 | 6.0 | 6.5 | 7.0 | 7.5 |
|-----------|-----|-----|-----|-----|-----|-----|-----|-----|
| 5.5       | 25.69 | 25.88 | 26.02 | 26.17 | 26.42 | 26.85 | 27.55 | 28.58 |
| 5.6       | 25.83 | 26.01 | 26.27 | 26.59 | 26.93 | 27.27 | 27.56 | 27.78 |
| 5.7       | 25.96 | 26.26 | 26.48 | 26.66 | 26.88 | 27.19 | 27.66 | 28.36 |
| 5.8       | 26.27 | 26.64 | 26.76 | 26.75 | 26.75 | 26.92 | 27.37 | 28.26 |
| 5.9       | 26.45 | 27.00 | 27.26 | 27.32 | 27.33 | 27.38 | 27.60 | 28.11 |
| 6.0       | 26.96 | 27.22 | 27.42 | 27.59 | 27.76 | 27.95 | 28.21 | 28.55 |
| 6.1       | 27.04 | 27.39 | 27.57 | 27.65 | 27.73 | 27.91 | 28.28 | 28.94 |
| 6.2       | 26.59 | 27.70 | 28.23 | 28.39 | 28.35 | 28.32 | 28.48 | 29.04 |
| 6.3       | 26.91 | 27.95 | 28.37 | 28.42 | 28.33 | 28.33 | 28.65 | 29.53 |
| 6.4       | 26.70 | 27.71 | 28.29 | 28.57 | 28.71 | 28.83 | 29.09 | 29.63 |
| 6.5       | 25.90 | 27.24 | 28.17 | 28.79 | 29.19 | 29.50 | 29.79 | 30.17 |
| 6.6       | 25.14 | 26.63 | 27.69 | 28.44 | 28.99 | 29.46 | 29.98 | 30.67 |
| 6.7       | 24.55 | 26.34 | 27.60 | 28.44 | 28.98 | 29.35 | 29.67 | 30.06 |
| 6.8       | 23.73 | 25.96 | 27.48 | 28.47 | 29.08 | 29.50 | 29.89 | 30.43 |
| 6.9       | 24.01 | 25.81 | 27.11 | 28.04 | 28.72 | 29.28 | 29.84 | 30.54 |
| 7.0       | 23.79 | 25.11 | 26.33 | 27.43 | 28.38 | 29.17 | 29.77 | 30.16 |
| 7.1       | 23.03 | 24.95 | 26.47 | 27.64 | 28.53 | 29.19 | 29.69 | 30.09 |
| 7.2       | 22.29 | 24.80 | 26.39 | 27.34 | 27.95 | 28.51 | 29.30 | 30.61 |
| 7.3       | 22.22 | 24.94 | 26.53 | 27.35 | 27.79 | 28.21 | 29.00 | 30.52 |
| 7.4       | 22.43 | 25.05 | 26.58 | 27.36 | 27.77 | 28.15 | 28.87 | 30.29 |
| 7.5       | 22.64 | 25.18 | 26.63 | 27.35 | 27.69 | 27.99 | 28.61 | 29.90 |
| 7.6       | 22.70 | 25.17 | 26.61 | 27.36 | 27.72 | 28.03 | 28.60 | 29.75 |
| 7.7       | 22.81 | 25.28 | 26.71 | 27.44 | 27.80 | 28.16 | 28.84 | 30.20 |
| 7.8       | 21.91 | 25.09 | 26.81 | 27.54 | 27.75 | 27.92 | 28.51 | 30.01 |

\( \log F_X \) (erg cm\(^{-2}\) s\(^{-1}\))

**Note.**

- Emission measures are log\(E_M\) values (units cm\(^{-5}\)) from Figure 10.
8. Exoplanet Effects on Coronal Abundances

Ever since the first detection of hot Jupiters orbiting very close to their host stars, there has been the question of what effects this close proximity will have on both the planet and star. There is no doubt that the planet characteristics will be greatly affected by being so close to the star. Its temperature close to their host stars, there has been the question of what activity. The WASP-18 system seems to be such a case, with cases where the exoplanet seems to be inhibiting stellar activity. The orbital period is about the same as the stellar rotation period, so this is a case where tidal synchronization may have taken place (Donati et al. 2008). Scharf et al. (2010) and Walker et al. (2008) find that the EM distribution are systematically discrepant from the EMs of other stars. Thus, we conclude that there is no evidence that the exoplanet is affecting the coronal temperature distribution, which looks normal for a star of this activity level. Coronal abundances are a different story, however. We have already noted that τ Boo A looks modestly discrepant in the relative abundance plots in Figures 7(a) and (b), and more convincingly discrepant in the absolute abundance plot in Figure 7(c).

This discrepancy was previously noted by Maggio et al. (2011) and Peretz et al. (2015) based on an XMM spectrum from 2003 June 24, although these authors were looking for evidence of changes in coronal abundance behavior induced by high photospheric metallicity, rather than changes that might be induced by the exoplanet of τ Boo A. One advantage of the Chandra spectrum analyzed here is that the τ Boo AB binary is resolved by Chandra, unlike XMM, meaning that our τ Boo A spectrum is not contaminated by any emission from the M2 V companion τ Boo B. This is explicitly demonstrated in Figure 12(a). Nevertheless, we find that the EM distribution and abundances that we measure for τ Boo A agree very well with those reported by Maggio et al. (2011), suggesting that τ Boo B is too faint to have significantly affected the analysis of the XMM data. In fact, direct comparison of the Chandra and
The XMM spectra shows only very subtle differences, demonstrating that Chandra/LTGS and XMM have impressively consistent flux calibrations, and also that τ Boo A does not vary much between the XMM observation in 2003 and the Chandra observation in 2017.

If planets can affect stellar coronae, it would be natural to assume that the effects might be greatest on the region of the star closest to the planet. Thus, coronal properties might be expected to vary with planetary orbital phase. The Chandra observations were fortuitously split into three pieces (see Table 2), which sample different orbital phases. Figure 12(b) shows the light curves of both τ Boo A and B measured from these separate exposures, based on the zeroth-order images. Using the ephemeris quoted by Catala et al. (2007) and Mengel et al. (2016), observation IDs 17715, 20019, and 20020 cover orbital phases $\phi = 0.69$–0.86, $\phi = 0.21$–0.31, and $\phi = 0.44$–0.49, respectively, and Figure 12(c) schematically shows the orbital geometry corresponding to these phases. The $\phi = 0$ phase corresponds to the first conjunction, when the planet is farthest behind the star from our perspective, with the planetary orbit tilted by $i = 44.5 \pm 1.5^\circ$ from the plane of the sky (Brogi et al. 2012).

The light curves in Figure 12(b) and the spectra of the three individual exposures reveal no clear variation with orbital phase. Furthermore, the 2003 XMM observations were taken at yet another distinct orbital phase of $\phi = 0.89$–1.13 (Maggio et al. 2011), and we have already noted the lack of significant variation between the XMM and Chandra/LTGS spectra.

We conclude that if the coronal abundance anomalies for τ Boo A apparent in Figure 7 are in fact due to the planet, the planet must be affecting the corona globally and not locally. This raises the question as to how an AR on τ Boo A would continue to be affected by a planet that has orbited behind the star. However, solar observations suggest that this may be possible in principle. In solar ARs, the FIP bias is observed to evolve on timescales comparable to the orbital timescale of τ Boo b. In particular, newly emerged solar ARs are generally found to have no FIP effect, and only acquire the normal solar FIP bias after several days (Widing & Feldman 2001). Therefore, one interpretation of the reduced FIP bias of τ Boo A is that the repetitive perturbations of the exoplanet,
with $P_{\text{orb}} = 3.31$ days, somehow reset the clock on the stellar ARs, making old ARs behave like young ones. Exactly how this might happen is a mystery. Are the gravitational or magnetospheric perturbations of the planet the most important?

Demonstrating that the exoplanet of $\tau$ Boo A is truly the cause of its coronal abundance anomalies requires finding examples of this effect for other exoplanet host stars. We take a first step in this direction using XMM measurements of O VIII and Fe XVII lines from another exoplanet host star, HD 189733A (K1.5 V), made by Pillitteri et al. (2014a). The transiting planetary companion of this star, HD 189733b, has an orbital period and semimajor axis of only $P_{\text{orb}} = 2.22$ days and $a = 0.031$ au, respectively, so this is another very close-in exoplanet system where tidal effects may be important. As mentioned above, HD 189733A is one of the two stars that Poppenhaeger & Wolk (2014) found to be more active than it should be compared to the activity level of its stellar companion, suggesting that the presence of HD 189733b may be leading to enhanced activity on HD 189733A. Using the density-sensitive O VII lines near 22 Å, Pillitteri et al. (2014a) report a potentially anomalous high coronal density of $n_e = (3-10) \times 10^{10}$ cm$^{-3}$ for HD 189733A, with 1σ error bounds. The best constraints on these lines for $\tau$ Boo A would be from the XMM measurements of Maggio et al. (2011) rather than our LETGS measurements. These line fluxes are not suggestive of high densities, although only a 1σ upper limit of $n_e < 3 \times 10^{10}$ cm$^{-3}$ can be quoted.

Wood et al. (2012) derived a relation between $F_{\text{bias}}$ and the Fe XVII/O VIII flux ratio using the Fe XVII lines at 15–17 Å and the O VIII line at 19.0 Å. Based on this relation and the line fluxes from Pillitteri et al. (2014a), we estimate $F_{\text{bias}} = 0.16$ for HD 189733A. However, we have to include the +0.084 correction factor discussed in Section 5, so our final value, with 1σ uncertainty, is $F_{\text{bias}} = 0.24 \pm 0.22$. In Figure 12(d), the FBST relation from Figure 7(b) is reproduced, and the locations of $\tau$ Boo A and HD 189733A are also shown. Both exoplanet host stars lie slightly above the main-sequence star relation. Extremely active stars such as EK Dra, AB Dor A, and AU Mic also lie above the relation (see Figure 7), but neither $\tau$ Boo A nor HD 189733A are in this high-activity regime, with X-ray luminosities of log $L_X = 28.76$ and log $L_X = 28.46$ (Schmitt & Liefer 2004), respectively. Other stars with these luminosities are consistent with the FBST relation. Thus, high activity probably does not account for the anomalous coronal abundances of $\tau$ Boo A or HD 189733A. Considering that previous observations have also suggested that stellar activity on these stars is being affected by their exoplanets, it seems reasonable to propose that the anomalously high $F_{\text{bias}}$ values are also somehow caused by the exoplanet influence.

9. Summary

We have conducted a survey of all main-sequence star Chandra/LETGS spectra with sufficient S/N for detailed analysis. This involves the analysis of 21 spectra from 19 stars, where both low-state and high-state spectra are considered for $\alpha$ Cen A and B. From these spectra we measure fluxes or upper limits for 118 lines from 49 ionic species. The EM analyses based on these measurements provide coronal temperature distributions and abundance measurements, leading to the following conclusions:

1. In contrast to past analyses, we no longer find narrow EM peaks at log $T = 6.6$, although there is still generally an EM maximum near that temperature, and we also find a systematic increase of +0.084 in the FIP bias parameter, $F_{\text{bias}}$, compared to past measurements. This is due to changes in Fe XVII line emissivities in version 7.1 of the CHIANTI database, used here as the source for atomic data. In particular, emissivities of three of the four strong Fe XVII lines at 15–17 Å increased by ~60% compared to past CHIANTI versions.

2. We expand on previous studies of the FBST relation, supplementing our LETGS sample with other published results from XMM and HETGS. Consideration of Altair (A7 V) and $\eta$ Lep (F1 V) allows us to extend the relation to earlier spectral types than before. We find that the relation is flat at A7-G5 spectral types, before increasing toward later types. Replacing spectral type with $T_{\text{eff}}$ allows us to derive a quantitative FBST relation, which is provided in Equation (5).

3. Absolute coronal Fe abundances are quantified using a line-to-continuum analysis, focusing on the 25–40 Å region. We find a roughly linear correlation between $[F_{\text{e}}/F_{\text{fe}}]$ and $T_{\text{eff}}$. $[F_{\text{e}}/F_{\text{fe}}] = -1.288 + 2.176 \times 10^{-4} T_{\text{eff}}$, albeit with a lot of scatter.

4. While the solar and stellar coronal abundances are perfectly consistent with each other with regard to relative abundances as quantified by $F_{\text{bias}}$, such is not the case for absolute abundances. While the solar FIP effect is generally accepted to involve low-FIP elements being enhanced, the stellar abundance measurements for solar-like G dwarfs seem more consistent with high-FIP depletions. This discrepancy casts doubt on the reliability of the line-to-continuum analysis.

5. We measure Ne/O = 0.39 for our sample of stars, consistent with previous work (Drake & Testa 2005). We find no indication of activity dependence that might mitigate the discrepancy with the solar value, which is nevertheless less dramatic than it used to be if the higher solar measurement from Young (2018), Ne/O = 0.24 ± 0.05, is used.

6. Coronal Si/Fe ratios are systematically high for M dwarfs. This may be due to the incorporation of Si into SiO in M-dwarf photospheres. We propose that the robustness of SiO leads to Si behaving more like a high-FIP element with regard to its coronal abundance.

7. Comparisons are made with the solar full-disk EM distributions from $\text{SDO}$/EVE (Schonfeld et al. 2017). The corona of $\alpha$ Cen B is the most solar-like in our sample, with a similar EM distribution and X-ray surface flux.

8. We find no change in coronal abundances between the low-state and high-state spectra of $\alpha$ Cen A and B, suggesting that coronal abundances do not vary with activity cycles.

9. Our sample of main-sequence stars provides a consistent picture for how EM distributions change with increasing $F_X$, with no evidence for any dependence on spectral type. Thus, we derive EM distributions as a function of $F_X$ (see Figure 10 and Table 7), which can be used as the basis for future theoretical studies of coronal heating, and can also be used to estimate EM distributions for any star with a measured broadband X-ray luminosity.
10. The stellar EM distributions are compared with solar QS, AR, and FL distributions. The QS distribution, with log $F_X = 3.61$, may represent a minimum emission level for main-sequence star coronae. Between log $F_X = 5.0$ and log $F_X = 6.5$, the stellar distributions agree very well with the solar AR distribution for log $T < 6.6$ in terms of both magnitude and slope, consistent with the idea that the surfaces of moderately active stars are completely filled with solar-like ARs. The sum of the AR and FL distributions represents a reasonable approximation of the stellar distribution we find at log $F_X \sim 6.0$. Completely covering the surface of a star with the X-class flares represented by our FL distribution leads to a decent approximation of the observed EM distribution of our most active stars, with log $F_X \sim 7.5$.

11. In deriving the solar FL distribution, we also inferred a relation between GOES X-class and X-ray luminosity: $\log L_X = 0.56 \log F_X + 27.75$.

12. The coronal abundances of the exoplanet host τ Boo A are anomalous. The $F_{\text{Fe}}$ value is somewhat higher than it should be, and $\log [Fe/Fe_{\odot}]$ is much lower. An $F_{\text{Fe}}$ estimate for another exoplanet host, HD 189733A, from XMM measurements is also somewhat higher than it should be. We therefore conclude that the close-in, massive exoplanets of these stars may be affecting their coronal abundances.

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Software: CIAO (v4.9; Fruscione et al. 2006), PINTofALE (v2.97; Kashyap & Drake 2000).
