X-ray Activity on the Star-Planet Interaction Candidate HD 179949

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ABSTRACT

We carry out detailed spectral and timing analyses of the Chandra X-ray data of HD179949, a prototypical example of a star with a close-in giant planet with possible star-planet interaction (SPI) effects. We find a low coronal abundance Fe/H≈0.2 relative to the solar photosphere, as well as lower abundances of high FIP elements O/Fe≲1, Ne/Fe≲0.1, but with indications of higher abundances of N and Al. This star also has an anomalous FIP bias of ≈0.03±0.03, larger than expected for stars of this type. We detect significant intensity variability over time scales ranging from 100 s - 10 ks, and also evidence for spectral variability over time scales of 1-10 ks. We combine the Chandra flux measurements with Swift and XMM-Newton measurements to detect periodicities, and determine that the dominant signal is tied to the stellar polar rotational period, consistent with expectations that the corona is rotational-pole dominated. We also find evidence for periodicity at both the planetary orbital frequency and at its beat frequency with the stellar polar rotational period, suggesting the presence of a magnetic connection between the planet and the stellar pole. If these periodicities represent a SPI signal, the lack of phase dependence in coronal temperature or flaring suggest that the SPI in this system is driven by a quasi-continuous form of heating (e.g., magnetic field stretching) rather than a highly sporadic, hot, impulsive form (e.g., flare-like reconnection).
1. INTRODUCTION

The first confirmed detection of an exoplanet occurred in 1992 (Wolszczan & Frail 1992), and since then, the number of exoplanet detections has been increasing rapidly, reaching ≈5000 confirmed detections by October 2022. Among them, giant planets are commonly detected, with ~450 such planets having semi-major axes < 0.15 AU and 0.5 ≤ M_{Jup} ≤ 15\(^1\).

(Ballesteros et al. 2019). Magnetic and tidal interactions between the planet and the host star in such systems can affect stellar activity (Rubenstein & Schaefer 2000; Cuntz et al. 2000). Such star-planet interactions (SPI) could have measurable effects (Shkolnik & Llama 2018; Strugarek 2021) on both chromospheric and X-ray output. For example, Shkolnik et al. (2003) identified chromospheric variability in HD 179949 tied to planetary phase and (Kashyap et al. 2008) claimed an average of 4× enhancement of X-ray luminosity in stars with close-in planets (though other studies have not confirmed this). Poppenhaeger et al. (2010) and Pillitteri et al. (2014a,b, 2022) have demonstrated that stellar activity in HD 189733 appears to be affected by the planetary phase and that a coeval star in the same system without a close-in giant planet is considerably weaker in activity.

HD 179949 is the prototypical example of a stellar system with a close-in giant planet which could be affected by SPI. It has a hot Jupiter orbiting it at a distance of ∼ 0.04 AU with a period of ∼3.1 days (Table 1). The star itself has an equatorial rotational period of 7.62 ± 0.05 days and a polar rotational period of 10.30 ± 0.80 days, (Fares et al. 2012). Cauley et al. (2019) found the star’s magnetic field strength to be 3.2±0.3 G.

Shkolnik et al. (2003) and Gurdemir et al. (2012) presented evidence for planet-related emission variability in HD 179949. This effect lasted for over a year and was correlated with the orbital phase of the planet, suggesting a magnetic interaction. Shkolnik et al. (2003) suggested that these interactions could produce a chromospheric hot spot which rotates in phase with the planet’s orbit, and is thus modulated by the orbital period. However, only intermittent variability rather than regular periodicity has been observed. Nevertheless, this intermittent variability is expected for some cases of magne-
Table 1. Stellar and Planetary parameters

| Property                           | Value                    | Reference                          |
|------------------------------------|--------------------------|------------------------------------|
| HD 179949: HIP/HIC 94645, HR 7291, GJ 749, Gumala |                          |                                    |
| Spectral Type                      | F8V                      | Houk & Smith-Moore (1988)           |
| Distance                           | 27.5 ± 0.6 pc            | Gaia Collaboration et al. (2018)    |
| Age                                | 1.20 ± 0.60 Gyr          | Bonfanti et al. (2016)              |
| Mass                               | 1.23 ± 0.01 M⊙          | Bonfanti et al. (2016)              |
| Radius                             | 1.20 ± 0.01 R⊙          | Bonfanti et al. (2016)              |
| mV, B − V                          | 6.237m, 0.534m          | Høg et al. (2000)                  |
| Proper motion (RA, Dec)            | (+118.52,-102.235) mas/yr| Gaia Collaboration et al. (2018)    |
| Equatorial rotational period       | 7.62 ± 0.05 days         | Fares et al. (2012)                 |
| Polar rotational period            | 10.3 ± 0.8 days          | Fares et al. (2012)                 |
| Magnetic field strength            | 3.2 ± 0.3 G              | Cauley et al. (2019)                |
| Bolometric Luminosity              | 1.95±0.01 L⊙            | Bonfanti et al. (2016)              |
| Photospheric [Fe/H]                | 0.056±0.055†            | Gonzalez & Laws (2007)              |
| Effective Temperature T_{eff}      | 6220 ± 28 K             | Bonfanti et al. (2016)              |
| Hot Jupiter HD 179949b             |                          |                                    |
| Mass                               | 0.980 ± 0.004 M_{Jupiter}| Butler et al. (2006)               |
| Radius                             | 1.05 R_{Jupiter}         | Wang & Ford (2011)                  |
| Orbital semi-major axis            | 0.044 ± 0.003 AU         | Butler et al. (2006)                |
| Orbital period                     | 3.09285 ± 0.00056 days   | Shkolnik et al. (2003)              |
| Ephemeris (φ = 0)                  | 2452479.823 ± 0.093 days| Shkolnik et al. (2003)              |

† Corrected to Anders & Grevesse (1989)

ranges spanning three consecutive orbits. We adopt the ephemeris of Shkolnik et al. (2003) (see Table 1) to describe the planetary orbit. Each observation covers ≈40° of the orbital phase, and the expected uncertainty in the absolute planetary phase at any given epoch is ≲25°. All the observations were carried out over one stellar polar rotation.

For each observation, we locate the source position by finding the centroid of photons inside a circle of radius 2″ centered at the proper motion corrected position. We extract the source counts from a circle of radius 3.6″ and background counts from an annulus with radii 10.8″ − 28.8″ corresponding to an area 55× larger than the source region. A summary of the observation specific source locations and strengths is in Table 3.

We perform detailed intensity and hardness ratio light curve analyses over several passbands. We choose both standard broad passbands (such as those defined by the Chandra Source Catalog, CSC) to allow comparisons with other analyses, as well as narrow passbands targeted to particular line complexes. The list of passbands is in Table 4. We typically merge the medium and hard bands due to the dearth of counts in the latter. The narrow bands are centered around strong emission lines in the H- and He-like ions of Oxygen, Neon, and Magnesium, and the Fe XVII-XVIII lines that dominate coronal radiation at ≈6 MK. Because the emission from these lines dominate at different temperatures, we expect that relative changes in intensity in the narrow passbands are diagnostic of temperature fluctuations as well as abundance variations. The count rates in each of the bands are shown in Figure 1.

To augment phase coverage, we also use fluxes from XMM-Newton Scandariato et al. (2013) and Swift D’Elia et al. (2013). The fluxes from XMM-Newton and Swift are shown in Figure 2 (along with Chandra fluxes estimated in Section 3.1). The Swift fluxes are those computed assuming a plasma temperature of 5 MK, and we discard three measurements from Swift when it appears to be undergoing flaring (with flux > 7×10^{-13} ergs cm\(^{-2}\) s\(^{-1}\)). The upper panel shows the fluxes as originally estimated, and the lower panel shows them rescaled such that the average of the fluxes in each of XMM-Newton and Swift match the Chandra average. We remove in this way the effects of possible long-term activity variations in periodicity estimations (see below).

3. ANALYSIS

3.1. Spectral Fits

We carry out a detailed analysis of the Chandra ACIS-S spectra of HD 179949 obtained during each observa-
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Table 2. Observation Log

| ObsID | Observation Start Time [UT] | Exposure [ksec] | Planetary phase range* [deg] |
|-------|-----------------------------|----------------|-------------------------------|
| 5427  | 2005-05-21 18:12:41         | 29.581         | [292.945 - 332.795] ± 24.303 |
| 6119  | 2005-05-22 11:37:35         | 29.648         | [17.490 - 57.431] ± 24.316 |
| 6120  | 2005-05-29 16:51:03         | 29.648         | [137.526 - 177.467] ± 24.453 |
| 6121  | 2005-05-30 12:00:30         | 29.644         | [203.933 - 243.870] ± 24.464 |
| 6122  | 2005-05-31 10:50:19         | 31.763         | [318.626 - (360+)1.416] ± 24.482 |

*Phase $\phi = 0$ is when HD 179949b is in conjunction, i.e., between HD 179949 and the observer along the line of sight. Error bars are computed by propagating the uncertainty in the phase to the observation epoch.

Table 3. Summary of Chandra observations

| ObsID | Measured Position (RA, Dec) | Offset from expected position† | Counts in source region* | Counts in background region* | Net Counts*‡ | Net Rate*‡ |
|-------|-----------------------------|------------------------------|--------------------------|----------------------------|--------------|-----------|
| 5427  | (19:15:33.2990, $-24^\circ$ 10' 46.127") | (+0.520",-0.011") | 1597                     | 124                       | 1590±41      | 54±1.4    |
| 6119  | (19:15:33.2853, $-24^\circ$ 10' 46.159") | (+0.305",-0.044") | 1300                     | 118                       | 1300±37      | 44±1.3    |
| 6120  | (19:15:33.2849, $-24^\circ$ 10' 46.084") | (+0.299",+0.033") | 1458                     | 117                       | 1460±39      | 49±1.3    |
| 6121  | (19:15:33.2850, $-24^\circ$ 10' 46.061") | (+0.304",+0.056") | 1599                     | 116                       | 1600±41      | 54±1.4    |
| 6122  | (19:15:33.2887, $-24^\circ$ 10' 46.049") | (+0.358",+0.069") | 1934                     | 497                       | 1920±45      | 61±1.4    |

† Based on positions from Gaia Collaboration et al. (2018), corrected for proper-motion.

‡ in broad band 0.5-7.0 keV.

Table 4. Passbands

| Band | Energy Range [keV] | Comment |
|------|-------------------|---------|
| u    | 0.2 - 0.5         | CSC ultrasoft |
| s    | 0.5 - 1.2         | CSC soft |
| m+h  | 1.2 - 8.0         | CSC medium + hard |
| b    | 0.5 - 7.0         | CSC broad |
|      | Line dominated bands |
| Oxy  | 0.5 - 0.7         | OVII & OVIII |
| Fe   | 0.7 - 0.9         | FeXVII and FeXVIII |
| Ne   | 0.9 - 1.2         | NeX and NeXI |
| Mg   | 1.2 - 3.0         | MgXIII and MgXII |

a collisionally excited thermal emission model for the corona (APEC; Brickhouse et al. 2000, Smith et al. 2001, Foster et al. 2016), and an interstellar absorption model (Tübingen-Boulder; see Wilms et al. 2000) with the H column density fixed at $N_\text{H} = 10^{19}$ cm$^{-2}$. In all cases we compute abundances relative to the solar photosphere as listed by Anders & Grevesse (1989) (with number abundances $\frac{A(C)}{A(H)} = 3.63 \times 10^{-4}$, $\frac{A(N)}{A(H)} = 1.12 \times 10^{-4}$, $\frac{A(O)}{A(H)} = 8.51 \times 10^{-4}$, $\frac{A(Fe)}{A(H)} = 1.23 \times 10^{-4}$, $\frac{A(Ne)}{A(H)} = 4.68 \times 10^{-5}$). We carry out the modeling in stages of increasing complexity, evaluating the quality of the fit at stage and continuing with more complex models only if necessary. Our fit results are thus designed to be robust against fluctuations, and describe the most information that may be reliably gleaned from the data. The models we consider are:
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1m An APEC model with one temperature component, fitting normalization, temperature, and metallicity;

2m An APEC model with two temperature components, fitting the normalization and temperature for each component separately and the metallicity jointly;

2v As in Model 2m, but allowing the abundances to vary freely in groups based on similar FIP (First Ionization Potential) values, as \{C,O,S\}, \{N,Ar\}, \{Ne\}, \{Mg,Al,Sn,Fe,Ni\}; and

2v/Z As in Model 2v, but the abundance of a selected element Z fitted separately. E.g., for model 2v/Al, the abundances are fit in groups \{C,O,S\}, \{N,Ar\}, \{Ne\}, \{Mg,Sn,Fe,Ni\}, \{Al\}. This is designed to explore whether enhancements of specific elements are measurable.

We carry out the fits in all cases by minimizing the c-statistic (cstat; see Cash 1979) over the passband 0.4-7.0 keV. Background is negligible (< 0.5%, see Table 3) in all cases and is ignored. We use the MCMC-based pyBLoCXS method (van Dyk et al. 2001) as implemented in Sherpa. We compute the observed cstat\text{obs}, and the nominal expected cstat\text{model}, and the variance \sigma^2\text{cstat} as described by Kaastra (2017), and reject the model as an adequate fit if the difference between the observed and model expected values, \Delta\text{cstat} = |cstat\text{obs} - cstat\text{model}| exceeds 2\sigma\text{cstat}.

When \Delta\text{cstat} is deemed acceptable, we further analyze the residuals to the spectral fit, since acceptable fits may still be accompanied by structures in residuals that indicate that some aspect of the model is inadequate over small energy ranges. We evaluate the quality of the residuals by computing the cumulative sum of the residuals (CuSum) and comparing it against a null distribution generated from the fitted model. For each iteration from pyBLoCXS, we generate a fake spectrum from the corresponding model parameters, and build up a set of CuSum curves. Then we calculate measures that test both the width and strength of the residual structure by carrying out two tests:

1. First, we compute the ±90% point-wise envelope of CuSum over the energy range 0.5-3.5 keV, and evaluate the percentage of bins, pct\text{CuSum}, where the observed CuSum exceeds the 90% bounds. If this percentage is \geq 10%, the model is considered inadequate and the next stage of complexity in the model is considered. If, on the other hand, the percentage is \leq 10% this is taken as a sign of overfitting, and the less-complex model is accepted.

2. Second, for each simulated spectrum, we calculate the total area of the CuSum curve that falls outside the 5-95% bounds from the ensemble of the simulations and construct a null distribution of the excess area beyond the 90% bounds. We then compute the same quantity for the observed CuSum curve, and hence the corresponding p-value (parea) with reference to the null distribution. This process is similar to the posterior predictive p-value calibration procedure described by Protassov et al. (2002). If this parea \leq 0.05, we consider the corresponding model an inadequate fit.

These tests are illustrated in Figure 3. The left panels in Figure 3 show two examples of pct\text{CuSum} calculated for two successive models (2m at the top and 2v at the bottom) for the same dataset (ObsID 6119); in both cases, the \Delta\text{cstat} is acceptable (+0.3 for model 2m and −0.86 for model 2v). Using just this information, there is no reason to prefer the more complex spectral
Figure 2. Measured X-ray fluxes of HD 179949 over phase Φ. The top panel shows raw fluxes obtained from XMM-Newton (blue points; Scandariato et al. 2013), Swift (green points; D’Elia et al. 2013), and the Chandra fluxes (red points) measured here (see Section 3.1). The bottom panel shows the same data, with the XMM-Newton and Swift fluxes separately normalized on average to Chandra. The grey bands show the Chandra phase coverage of HD 179949b.

We report the results of the spectral fits for each ObsID in Table 5 for the best acceptable model obtained for each case, along with the cstat, Δcstat, pctCuSum, and parea for the accepted model. The corresponding fits and residuals are shown in Figure 4, and the CuSum plots and distributions are shown in Figure 12 in Appendix B. We note that the data are never fit well with model 1m, and never require models more complex than model 2v. ObsID 6121 is fit adequately with model 2m; model 2v can be rejected due to an abnormally low pctCuSum = 1.10%. In no case is model 2v/Z or other complex modifications statistically justifiable.
While some exploratory fits indicated that allowing Al to be freely fit produced better fit statistics, there is no reason to choose those fits over those of model 2v. In addition to the plasma temperatures, component normalizations, abundances, and the corresponding fluxes, we also compute the FIP bias \( F_{\text{bias}} \) (Wood et al. 2012, Testa et al. 2015, Wood et al. 2018). The FIP bias is a summary of how much the abundances of high-FIP elements like C, N, O, and Ne deviate relative to a low-FIP element like Fe, in the corona relative to the photosphere, normalized to solar abundances,

\[
F_{\text{bias}} = \frac{1}{4} \sum_{X=C,N,O,Ne} \log \frac{X/Fe}_{\text{cor}} - \log \frac{X/Fe}_{\text{phot}}.
\]

(1)

3.2. Timing Analysis

3.2.1. Light Curves

In order to explore the temporal variability in HD 179949, we construct light curves of counts in all the passbands of interest (see Table 4). We employ the Bayesian Blocks algorithm (Scargle et al. 2013) to build adaptively binned count rates, using the implementation by Astropy Collaboration et al. (2022).

We modify the process by which the blocks are constructed in two ways: first, because the CCD frame readout time is \( \tau = 3.24104 \) s, we discard any change points that occur < 10\( \tau \) after another; second, we recalibrate the parameter \( p_0 \) that controls the false alarm probability (FAP) to both account for the 10\( \tau \) filter as well as to calibrate the steepness of the prior distribution. We accomplish this via simulations involving the same number of photons as in the observed data, 9963, spaced uniformly over a normalized time duration equal to the full observation duration, \( 152295/\tau \). That is, each simulation represents events that would be obtained from a flat light curve that has zero true change points. Change points for several values of \( p_0 \) were computed for the same set of event times. We discard all change points that follow another by < 10, and record the number of change points \( n_{\text{sim}}^{cp}(p_0) \). Because the simulated events...
have zero true change points, every change point found, at any given \( p_0 \), represents false positives. The fraction

\[
p_{\text{true}}(p_0) = \frac{n_{\text{CP}}(p_0)}{N_{\text{sim}}}
\]

thus represents the actual false alarm probability parameterized by \( p_0 \) for data sets of the size we are dealing with. We repeat the simulations 50 times, and estimate \( p_{\text{true}}(p_0) \) as the average over the simulations at each given \( p_0 \). The result of these computations is shown in Figure 3.2.1. The range of \( p_0 \in (0, 0.55) \) is approximately linear, and we fit a quartic polynomial to this range and extrapolate it to smaller values of \( p_{\text{true}} \). We find that a value of \( p_{\text{true}} = 0.01 \), corresponding to a 1% false alarm probability, is obtained for \( p_0 \approx 0.06 \), and hence adopt \( p_0 = 0.06 \) for all the change point calculations carried out below.
Figure 4. Model fits (Table 5) and residuals for each Chandra observation. The data are shown as the blue points with Gehrels error bars, grouped by 5 counts in each bin for the sake of visualization, and the model is shown as the orange stepped curve. The residuals relative to the errors are shown in the lower segment of each panel.

Figure 5. Calibrating the correction factor for the False Alarm Probability (FAP; $p_{true}$) for the AstroPy implementation of the Bayesian Blocks algorithm. We fit the $p_0$ value obtained from simulations to the input $p_0$ value using a quartic best-fit equation, after merging all change points separated by $<$30 sec. This provides a corrected $p_0 = 6\%$ as the appropriate value to use for a 1% FAP. We compute the block representation of light curves using this value (see, e.g., Figure 6).

Table 6. Summarizing the variability in light curves

| ObsID | Parameters | 5427 | 6119 | 6120 | 6121 | 6122 |
|-------|------------|------|------|------|------|------|
| Oxy   | $\langle \Delta t_{\text{bin}} \rangle$ | 257\tau | 193\tau | 226\tau | 309\tau | 237\tau |
|       | $N_{\text{blocks}}$               | 36   | 48   | 41   | 30   | 42   |
|       | $\chi^2_{\text{red}}$             | 15.3 | 7.7  | 14.0 | 5.1  | 9.4  |
| Fe    | $\langle \Delta t_{\text{bin}} \rangle$ | 264\tau | 206\tau | 238\tau | 221\tau | 331\tau |
|       | $N_{\text{blocks}}$               | 35   | 45   | 39   | 42   | 32   |
|       | $\chi^2_{\text{red}}$             | 8.1  | 11.1 | 7.9  | 7.9  | 8.4  |
| Ne    | $\langle \Delta t_{\text{bin}} \rangle$ | 250\tau | 215\tau | 215\tau | 309\tau | 255\tau |
|       | $N_{\text{blocks}}$               | 37   | 43   | 43   | 30   | 39   |
|       | $\chi^2_{\text{red}}$             | 16.1 | 13.7 | 16.3 | 6.7  | 6.2  |
| Mg    | $\langle \Delta t_{\text{bin}} \rangle$ | 257\tau | 201\tau | 281\tau | 238\tau | 267\tau |
|       | $N_{\text{blocks}}$               | 36   | 46   | 33   | 39   | 37   |
|       | $\chi^2_{\text{red}}$             | 12.5 | 11.4 | 9.3  | 15.9 | 8.1  |
| u     | $\langle \Delta t_{\text{bin}} \rangle$ | 237\tau | 265\tau | 290\tau | 226\tau | 231\tau |
|       | $N_{\text{blocks}}$               | 39   | 35   | 32   | 41   | 43   |
|       | $\chi^2_{\text{red}}$             | 7.6  | 9.3  | 7.9  | 14.8 | 8.8  |
| s     | $\langle \Delta t_{\text{bin}} \rangle$ | 201\tau | 226\tau | 265\tau | 244\tau | 343\tau |
|       | $N_{\text{blocks}}$               | 46   | 41   | 35   | 38   | 29   |
|       | $\chi^2_{\text{red}}$             | 6.9  | 9.4  | 10.0 | 8.7  | 11.7 |
| m+h   | $\langle \Delta t_{\text{bin}} \rangle$ | 257\tau | 201\tau | 273\tau | 238\tau | 255\tau |
|       | $N_{\text{blocks}}$               | 36   | 46   | 34   | 39   | 39   |
|       | $\chi^2_{\text{red}}$             | 12.3 | 11.7 | 9.6  | 15.9 | 8.7  |
| broad | $\langle \Delta t_{\text{bin}} \rangle$ | 250\tau | 211\tau | 299\tau | 265\tau | 293\tau |
|       | $N_{\text{blocks}}$               | 37   | 44   | 31   | 35   | 34   |
|       | $\chi^2_{\text{red}}$             | 6.9  | 10.8 | 8.9  | 5.9  | 7.7  |
We show the Bayesian Blocks count rate light curves for the broad band in Figure 6 for all five datasets. We compute several measures of variability:

**Average Block size:** $\langle \Delta t_{\text{bin}} \rangle$ is the average of the block sizes $\Delta t_{\text{bin}}$ found with Bayesian Blocks, and represents the time scale over which the algorithm requires changes in its piecewise constant light curve model.

**Number of blocks:** The number of blocks, $N_{\text{blocks}} = n^CP - 1$, defined by $n^CP$ change points found by Bayesian Blocks. This number complements $\langle \Delta t_{\text{bin}} \rangle$ to show the range and ubiquity of variability.

**Chi-square:** A reduced $\chi^2$, defined as

$$\chi^2_{\text{red}} = \frac{1}{(N_{\text{blocks}} - 1)} \sum_B \frac{(\text{rate}_B - \text{rate}_{\text{ObsID}})^2}{\sigma_B^2},$$

where $\text{rate}_B = \frac{\text{counts}_B}{\Delta t_B}$ is the observed rate in block $B$ when $\text{counts}_B$ counts are observed over a duration $\Delta t_B$, $\text{rate}_{\text{ObsID}}$ is the count rate for the full dataset (the orange horizontal line in Figure 6), and $\sigma_B = \sqrt{\frac{\text{counts}_B}{\Delta t_B}}$ is the uncertainty on the rate. This represents the quality with which a model with unvarying intensity can be fit to the light curves.

These quantities are reported in Table 6 for all the passbands and for all ObsIDs, and show that there is clear evidence for variability in the light curves over a broad range of time scales from $\approx 100$ s to $\sim 5$ ks. We confirm that variability in HD 179949 is ubiquitous across passbands and datasets, with $\chi^2_{\text{red}}$ is invariably significantly $\gg 1$. The light curves for the remaining passbands are available via Zenodo\(^2\).

### 3.2.2. Hardness ratios

There are insufficient counts in the blocks found for the counts light curves (Section 3.2.1) to allow time-resolved spectroscopy. Instead, we calculate the evolution of hardness ratios over time and evaluate the strength of the evidence for spectral variability (for analyses that highlight how hardness ratios can be useful, see, e.g., Noel et al. 2022; Di Stefano et al. 2021). Specifically, we compute the color

$$C[A/B] = \log_{10} \frac{\text{counts in band A}}{\text{counts in band B}},$$

for various combinations of pairs of passbands in Table 4. We consider different combinations of the CSC bands and the line dominated bands. Variations in

\(^2\) doi:10.5281/zenodo.7220014; Acharya et al. (2022)
plasma temperatures would be reflected in changes in the CSC colors, while abundance variations could become discernible in the line dominated band ratios. We use BEHR (Bayesian Estimation of Hardness Ratios; Park et al. 2006) to account for background and estimate the most probable values (modes) and uncertainties (68% highest posterior density [HPD] intervals) of the color in each time interval considered.

In order to define appropriate time intervals, we use the change points found in the light curves by Bayesian Blocks (Section 3.2.1) analysis of the corresponding passbands. These change points are then combined and pruned to form time bins, and the color is computed independently for counts collected in each bin. The process is described in detail below.

1. We begin by constructing the union of the set of change points from both passbands to form a single set of change points in temporally sorted order.

2. All pairs of change points which are separated by < 10σ≈32 s are then grouped together and replaced by their average, for consistency with how the change points are initially determined. This grouping and averaging is carried out iteratively until no such pairs are found. We call these merged change points close mergers.

3. We repeat the above pair-wise merging process for all remaining points which are separated by < 30σ≈100 s. While this removes any sensitivity in our process to spectral changes at time scales shorter than 100 s, the likelihood of such quick changes being visible in low-density hot coronal plasma dominated by radiative cooling processes is low. We name such merged points loose mergers.

The time points left in the set at the end of the above merging procedure are used as the bin boundaries for events collected in both passbands. We show an example of variations in the color ratio of the Oxy and Fe passbands in Figure 7. The figure shows the mode of C[Oxy/Fe] (blue stepped curve) and associated 68% HPD uncertainty (blue vertical bars), along with the close mergers (red vertical lines) and loose mergers (grey vertical lines). The color estimate obtained for the full dataset is also shown (orange horizontal line), illustrating the scope of departure from constancy. As with the count rate light curves, we again compute a reduced $\chi^2$ as a measure of variability,

\[
\chi^2_{red} = \frac{1}{N_{bins} - 1} \sum_k \frac{\left( \langle C[A/B]_k \rangle - \langle C[A/B]_{ObsID} \rangle \right)^2}{\text{Var}[C[A/B]_k]},
\]

where $N_{bins}$ are the number of bins in the light curve of $C[A/B]$ after the merging process, $\langle C[A/B]_k \rangle$ is the mean color estimated from BEHR draws in each bin $k$, Var$[C[A/B]_k]$ is the corresponding variance, and $\langle C[A/B]_{ObsID} \rangle$ is the mean color estimated for the full ObsID. These $\chi^2_{red}$ values are reported in Table 8, with values of $\chi^2_{red} > 1 + \sqrt{\frac{2}{N_{bins}-1}}$ marked in italic, and $\chi^2_{red} > 1 + 3\sqrt{\frac{2}{N_{bins}-1}}$ marked in bold. It is apparent that the magnitude of the spectral variations is significantly smaller than the intensity variations. This suggests that count rates in the different bands tend to rise and fall in tandem, with most of the variations attributable to emission measure changes. Nevertheless, spectral variations over time scales of the observation duration is definitely present and are indeed detectable, as seen in several cases where significant deviations in $\chi^2_{red}$ is seen. Furthermore, detailed examination of the color light curves can yield information on smaller time

| ObsID | 5427 | 6119 | 6120 | 6121 | 6122 |
|-------|------|------|------|------|------|
| C[u/s] | $-0.52^{+0.03}_{-0.01}$ | $-0.48^{+0.03}_{-0.02}$ | $-0.48^{+0.03}_{-0.04}$ | $-0.52^{+0.03}_{-0.02}$ | $-0.59^{+0.03}_{-0.01}$ |
| C[u/m+h] | $+0.40^{+0.03}_{-0.04}$ | $+0.53^{+0.06}_{-0.04}$ | $+0.46^{+0.04}_{-0.04}$ | $+0.43^{+0.03}_{-0.04}$ | $+0.30^{+0.03}_{-0.04}$ |
| C[s/m+h] | $+0.90^{+0.04}_{-0.03}$ | $+1.02^{+0.05}_{-0.05}$ | $+0.94^{+0.03}_{-0.04}$ | $+0.94^{+0.03}_{-0.04}$ | $+0.87^{+0.02}_{-0.04}$ |
| C[Oxy/Fe] | $-0.40^{+0.03}_{-0.04}$ | $-0.37^{+0.03}_{-0.03}$ | $-0.36^{+0.03}_{-0.03}$ | $-0.38^{+0.03}_{-0.03}$ | $-0.39^{+0.04}_{-0.02}$ |
| C[Oxy/Ne] | $-0.03^{+0.04}_{-0.03}$ | $+0.11^{+0.04}_{-0.04}$ | $+0.04^{+0.03}_{-0.04}$ | $+0.04^{+0.04}_{-0.04}$ | $-0.06^{+0.03}_{-0.03}$ |
| C[Oxy/Mg] | $+0.25^{+0.04}_{-0.05}$ | $+0.39^{+0.04}_{-0.05}$ | $+0.31^{+0.04}_{-0.05}$ | $+0.29^{+0.04}_{-0.04}$ | $+0.21^{+0.03}_{-0.03}$ |
| C[Fe/Ne] | $+0.37^{+0.03}_{-0.04}$ | $+0.47^{+0.03}_{-0.04}$ | $+0.39^{+0.02}_{-0.02}$ | $+0.34^{+0.02}_{-0.02}$ | $+0.32^{+0.02}_{-0.02}$ |
| C[Fe/Mg] | $+0.65^{+0.04}_{-0.05}$ | $+0.75^{+0.05}_{-0.04}$ | $+0.67^{+0.04}_{-0.04}$ | $+0.67^{+0.03}_{-0.04}$ | $+0.59^{+0.03}_{-0.03}$ |
| C[Ne/Mg] | $+0.27^{+0.04}_{-0.06}$ | $+0.30^{+0.04}_{-0.06}$ | $+0.28^{+0.03}_{-0.05}$ | $+0.33^{+0.03}_{-0.05}$ | $+0.26^{+0.03}_{-0.03}$ |

* Showing the mode and the 68% highest posterior density intervals.
scale variations, as, e.g., the large rise in \(C\{\text{Oxy/Fe}\}\) seen in ObsID 5427 between times 11-15 ks after the start of the observation (top panel of Figure 7). A similar, but smaller variation is also seen in \(C\{\text{Oxy/Ne}\}\), but no variability is present in either the CSC band colors \(C\{s/m + h\}\) or in \(C\{\text{Fe/Ne}\}\). This is a strong indication of a temporary surge in the abundance of Oxygen, reminiscent of variations seen in streamers and plumes around active regions on the Sun (see, e.g., Raymond et al. 1998; Raymond 1999; Guennou et al. 2015). All of the color \(C\{A/B\}\) spectral variability plots are available via Zenodo\(^2\).

### Table 8. Summary of spectral variability\(^*\)

| ObsID  | 5427 | 6119 | 6120 | 6121 | 6122 |
|--------|------|------|------|------|------|
| \(C[u/s]\) | 1.4 (59) | 1.6 (56) | 1.2 (48) | 1.9 (55) | 1.8 (54) |
| \(C[u/m-h]\) | 0.9 (44) | 0.8 (39) | 1.2 (38) | 1.1 (43) | 1.1 (48) |
| \(C[s/m+h]\) | 1.6 (49) | 0.9 (41) | 1.3 (46) | 1.8 (45) | 0.8 (41) |
| \(C\{\text{Oxy/Fe}\}\) | 1.1 (38) | 1.6 (52) | 1.1 (47) | 1.0 (50) | 1.7 (45) |
| \(C\{\text{Oxy/Ne}\}\) | 1.3 (39) | 1.3 (57) | 1.2 (46) | 1.1 (38) | 1.5 (51) |
| \(C\{\text{Oxy/Mg}\}\) | 1.2 (35) | 0.6 (34) | 1.0 (40) | 0.9 (33) | 0.8 (44) |
| \(C\{\text{Fe/Ne}\}\) | 1.0 (46) | 1.1 (56) | 1.5 (51) | 1.0 (51) | 1.4 (51) |
| \(C\{\text{Fe/Mg}\}\) | 1.4 (43) | 1.1 (41) | 1.2 (46) | 1.2 (45) | 0.8 (40) |
| \(C\{\text{Ne/Mg}\}\) | 1.0 (41) | 0.9 (35) | 1.2 (42) | 1.3 (38) | 0.9 (49) |

\(^*\)The values of \(\chi^2_{red}\) (Equation 4) are shown, along with the number of bins \(N_{bins}\) in parentheses. Values with deviations > 1σ are marked in italic and deviations > 3σ are marked in bold.

#### 3.2.3. Lomb-Scargle Periodogram

After fitting the spectral models for the 5 datasets, we calculate the resultant fluxes in the 0.15—4.0 keV energy range. The results are shown in Table 5 and in Figure 2 along with flux data from XMM-Newton (Scandariato et al. 2013) and Swift (D’Elia et al. 2013). We choose to compare against Swift fluxes obtained assuming a plasma with temperature \(\approx\)5 MK as that is comparable to the plasma temperatures we find (Table 5. We remove three flaring events (with flux \(\gg7\times10^{-13}\) ergs s cm\(^{-2}\)) from the Swift dataset to focus only on periodic variability of the quiescent flux. In addition, we normalise the XMM-Newton and Swift flux values to the average \(\text{Chandra}\) flux to remove systematics like calibration uncertainties and long-term activity variations. The resultant fluxes are shown in the bottom plot of Figure 2.

We search for periodicity in the combined flux data \(f\) using the Lomb-Scargle (L-S) periodogram \(LS(f)\) (Lomb 1976; Scargle 1982), as implemented by VanderPlas (2018) (see left panel of Figure 9). In order to

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**Figure 7.** Color Hardness Ratios for Oxy and Fe bands \(C\{\text{Oxy/Fe}\}\) for the 5 datasets. The most probable value in each time bin (blue stepped curve) and the 68% HPD interval (blue vertical bars) are shown, along with the mode value measured for the full dataset (orange horizontal bar; the width represents the 68% HPD interval). The locations of close merger change points (red vertical lines) and loose merger change points (grey vertical lines) are also marked.
obtain better resolution in phase, we split the Chandra data further into segments of \( \geq 5 \) ks in each ObsID, converting the net counts to fluxes using a counts-to-energy conversion factor computed separately for each ObsID as the ratio of the flux measured (Table 5) to the net count rate (Table 3) for that ObsID.

We account for the effects of statistical fluctuations and the windowing function via bootstrapping. We first scramble the fluxes \( f \) using a random permutation operator \( Scr(\cdot) \) to obtain a new set of fluxes \( g_k = Scr_k(f) \), where \( k = 1, \ldots, 5000 \). We compute \( LS(g_k) \) for each of these permutations. When averaged, \( \langle LS(g) \rangle \) represents a periodogram that is devoid (due to the scrambling) of an intrinsic periodic signal, but includes effects due to spacing and data level. The scatter in the bootstrapped periodograms, \( \text{stddev}[LS(g)] \) provides a measure of the statistical and data spacing noise. The difference \( LS(f) - \langle LS(g) \rangle \) represents the residual signal that can be attributed to intrinsic periodicity.

We then compute the window function by constructing the L-S periodograms of a unit flux dataset sampled at the same times as \( f \). In order to ameliorate numerical instabilities, we allow for a jitter in the unit flux, obtaining values in each case drawn from a Gaussian with mean 1 and width equal to the fractional error on each flux value. We calculate the window function 5000 times and average the resulting periodograms to obtain \( \langle LS(W) \rangle \) (see right panel of Figure 9).

We then extract the de-aliased, window-scaled, signal of intrinsic periodicity

\[
P(f) = \frac{LS(f) - \langle LS(g) \rangle}{\langle LS(W) \rangle}.
\]

This scaled periodogram is shown in Figure 10 as the magenta curve, with select periods marked with vertical bars. We also show the expected 2\( \sigma \) uncertainty, constructed as \( 2 \times \text{stddev}[LS(W)] \), and shown as the grey shaded region (for visibility, the 2\( \sigma \) curve is placed above \( P(f) \), covering it where it crosses. It is clear that several peaks are present in \( P(f) \) at or close to periods of relevance to the HD 179949 system: we detect periodicities at the stellar polar rotational period (\( \sim 10 \) d), the planetary orbital period (\( \sim 3 \) d), and the beat period of the stellar polar and planetary orbital periods (\( \sim 4.5 \) d, consistent with some Ca II H&K modulations found by Fares et al. (2012)) at significances > 95%. While each of these are not definitively above the usually adopted "3\( \sigma \) threshold, the confluence of all three inter-related signals suggests that these periodicities do exist in the HD 179949 system.

4. DISCUSSION

4.1. Stellar Context

We further explore the properties of HD 179949 in the context of similar stars (see Table 9), by comparing the relative X-ray luminosity \( \frac{L_x}{L_{bol}} \), effective temperature, rotation period and metallicity. In Figure 11, we show the variation in \( \frac{L_x}{L_{bol}} \) with the Rossby number \( R_0 \) (the ratio of the equatorial rotational period \( P_{rot} \) and the convective turnover time scale \( \tau_c \)), where HD 179949 is represented with a red block whose height represents the variation in measured \( L_X \), and for \( R_0 \) calculated using \( \tau_c \) from Gunn et al. (1998). HD 179949 is unremarkable in this space, with a slightly lower activity than expected by the trend line, but consistent with other F dwarfs of similar milieu. Other physical parameters like effective temperature are also consistent with other F7V-F9V stars.
Figure 9. Left: The Lomb-Scargle periodogram $LS(f)$ for the normalized fluxes data $f$ from Chandra, XMM-Newton, and Swift. Right: The averaged window function $⟨LS(W)⟩$.

Figure 10. Analysis of periodicities in HD 179949. $P(f) = \frac{LS(f) - ⟨LS(g)⟩}{⟨LS(W)⟩}$ which is the modified data periodogram, scaled by the averaged window function (magenta) is compared with $2 \cdot \text{stddev}[LS(g)]/⟨LS(W)⟩$ which is twice the noise in the averaged time-scrambled periodogram similarly normalized (grey). Peaks with SNR $> 2$ are considered possible periodicities of flux variation. Here, a clear sign of variability tied to the stellar polar rotation ($f_{\text{Polar}}$) can be seen. An argument for possible periodicities tied to the planetary orbital frequency ($f_{\text{Orbital}}$), and the beat frequency ($f_{\text{beat}}$) between $f_{\text{Polar}}$ and $f_{\text{Orbital}}$ can also be made. Periodicity tied to the stellar equatorial rotation ($f_{\text{Equatorial}}$) is not detected.
In contrast, we find that the photospheric abundances are higher than similar stars. Typically, $[\text{Fe}/\text{H}]_{\text{ph}} \lesssim -0.1$ for dF stars of similar activity and distance, but is $\gtrsim 0$ for HD 179949 (Gonzalez & Laws 2007; corrected to Anders & Grevesse 1989). Note however that this appears to be a characteristic of the presence of close-in giant planets (Fischer & Valenti 2005; Wang & Fischer 2015); indeed, for the stars listed in Table 9, we find that for those without confirmed exoplanets (14 stars), $\langle [\text{Fe}/\text{H}]_{\text{ph}} \rangle = -0.22 \pm 0.18$ while for stars with confirmed exoplanets (3, including HD 179949), $\langle [\text{Fe}/\text{H}]_{\text{ph}} \rangle = -0.06 \pm 0.10$.

4.2. Coronal Abundances

We find consistent coronal abundance measurements for O/Fe and Ne/Fe across all datasets, with $\text{O/Fe} \sim 0.7$ and $\text{Ne/Fe} \ll 0.1$. In contrast, $\text{N/Fe}$ is found to be $\gtrsim 1.0$ for some datasets (ObsIDs 5427, 6119, and 6120). Further, the average $Z_{\text{Fe}} \sim 0.2$ for all datasets except 6119, where it rises to $\approx 0.45$. Note that for ObsID 6119 (phase $\phi$ occurring just after planetary conjunction), the abundances of other metals are boosted as well, though within the uncertainty bounds from other ObsIDs. Overall, we conclude that the corona of HD 179949 is significantly deficient in Fe relative to the Sun. We also find indications of abundance variations at small time scales in the color light curves. For instance, we find an extended duration ($\gtrsim 3$ ks) where $C[\text{Oxy}/\text{Fe}]$ and $C[\text{Oxy}/\text{Ne}]$ increase (see Section 3.2.2) without an accompanying change in the temperature sensitive $C[\text{s/m + h}]$ or in $C[\text{Fe}/\text{Ne}]$, suggesting a short duration surge in O abundance. However, detailed spectral analyses, exploring whether the abundances in specific elements can be estimated, are statistically unsupported. None of the models $2\nu/2$ were found to be required to explain the spectra, though some instances of models of $2\nu/\text{Al}$ suggest that Al/Fe $> \text{solar}$.

We calculate the mean and standard deviations for the coronal $[\text{C}/\text{Fe}]$, $[\text{O}/\text{Fe}]$, $[\text{N}/\text{Fe}]$ and $[\text{Ne}/\text{Fe}]$ abundance ratios from the posterior distributions of abundances obtained during the spectral fits (Table 5). Abundance ratios for photospheric C, O and N were taken from Luck (2018), corrected to a baseline solar abundance of (Anders & Grevesse 1989), as $[\text{C}/\text{Fe}]_{\text{ph}} = -0.06$, $[\text{N}/\text{Fe}]_{\text{ph}} = 0.12$, $[\text{O}/\text{Fe}]_{\text{ph}} = 0.12$. We also adopt an error on these abundance ratios equal to the reported $\sigma_{[\text{Fe}]/[\text{ph}]} = 0.06$. For Ne, the value of $[\text{Ne}/\text{O}]_{\text{ph}} \sim -0.39$ from Drake & Testa (2005). These are then used to calculate the $F_{\text{bias}}$ following Wood et al. (2012), as a way to characterize the FIP effect. As shown in Wood et al., $F_{\text{bias}} < 0$ indicates solar-like FIP effect, and $F_{\text{bias}} > 0$ indicates an inverse FIP-effect. Wood et al. found a trend of decreasing $F_{\text{bias}}$ going from dM to dG. This trend is explored up to F-type stars in Wood et al. (2018), to conclude that $F_{\text{bias}}$ of most F-type stars should be $\lesssim -0.5$, which agreed with Seli et al. (2022), who reported $F_{\text{bias}} = -0.53 \pm 0.08$ for HD 179949. However, we measure $F_{\text{bias}}$ ranging from $-0.08$ to $0.18$ in the 5 Chandra datasets, with a standard deviation of $\sim 0.50$ dex (Table 5). This makes HD 179949 similar to the exceptional cases of $\tau$ Boo A and HD 189733A identified in Wood et al. (2018) which have a higher $F_{\text{bias}}$ than expected and they attributed this to the possible influence of the close-in Jupiter mass planets. A similar scenario could explain the anomalously high $F_{\text{bias}}$ for HD 179949, resulting in a weak inverse FIP effect instead of the expected solar-like behaviour. Testa et al. (2015) noted that low-activity stars usually show solar-like FIP effects, further pointing out to the anomalous $F_{\text{bias}}$ value of HD 179949.

4.3. Spectral Variations

As we show in Section 3.2, considerable variability is present in both intensity and color over a range of time scales. We first note that the existence of close and loose mergers of change points from several bands point to there being substantial energy release incidents that reinforce the presence of ubiquitous variability in the corona of HD 179949. Several passband color combinations show the presence of variability over the durations.
and observations; we find that all passband combination for all passband combinations mated goodness of a constant hardness ratio model for observations (see Section 3.2.2). Table 8 lists the esti-
ti light curves. We find that pervasive variability at the presence of spectral variability through hardness ra-
to carry out time-resolved spectroscopy, we can infer changes (see Figures 7,8).

This variability suggests that the corona of HD 179949 is dynamic, with intermittent impulsive energy releases occurring continuously. This small time scale variability also translates to the corona in its gross characteristics over the ~8-hour observation durations. Variations of factors of 2× is present in all fitted parameters (Table 5). Nevertheless, it is notable that large flares are not present. We speculate that this is due to the inhibition of large stresses from being built up in the magnetic loops in the corona because periodic interference of the planetary magnetic field which would act to dissipate the stresses continuously. The estimated F_{\text{bias}} value shows that high FIP elements are depleted in HD 179949. Further, due to the under-abundance of Neon, following from Laming (2015), we conclude that the star is not active. A comparison of $L_X/L_{\text{bol}}$ which is found to be [-5.6, -5.5] for HD 179949 with other F-type stars like HD 17156 with $L_X/L_{\text{bol}} = [-5.8, -6.0]$ further supports the state of quiescence (Maggio et al. 2015).

### 4.4. Magnetic SPI

The star’s large-scale magnetic geometry which has been mapped using ZDI (Semel 1989; Donati & Brown 1997) and near contemporaneously with Chandra observations during 2007 and 2009 (Fares et al. 2012). The well-observed 2009 data yield a strongly dipolar configuration, but with the magnetic pole inclined 70° to the rotational axis. Our detection of periodicity corresponding to the polar rotational period thus implies magnetic activity around the magnetic equator but confined to

| Name          | Spectral type | B - V  | $T_{\text{eff}}$ [K] | [Fe/H]_{ph} | log($L_X/L_{\text{bol}}$) | $R_0$ | Reference |
|---------------|---------------|--------|----------------------|-------------|--------------------------|-------|-----------|
| β Vir         | F8IV-F8V      | 0.550  | 6071                 | +0.03$^{+0.09}_{-0.07}$ | -5.65        | 0.628 | (1,2,3,4,11) |
| θ Per         | F7V           | 0.514  | 6196                 | -0.16$^{+0.07}_{-0.11}$ | -5.81        | 0.621 | (1,2,3,4,11) |
| ζ dor         | F7V-F8V       | 0.526  | 6153                 | -0.40$^{+0.09}_{-0.07}$ | -4.60        | 2.913 | (1,2,3,4,11) |
| 36 UMa        | F8V           | 0.515  | 6172                 | -0.25$^{+0.09}_{-0.11}$ | -5.43        | 0.724 | (1,2,3,4,11) |
| ν And†        | F8V           | 0.540  | 6106                 | -0.07$^{+0.09}_{-0.07}$ | -5.85        | 0.621 | (1,2,4,5,11) |
| ε Psc         | F7V           | 0.500  | 6241                 | -0.24$^{+0.09}_{-0.11}$ | -5.92        | 0.628 | (1,2,3,4,11) |
| 99 Her        | F7V           | 0.520  | 5947                 | -0.73$^{+0.09}_{-0.11}$ | -5.83        | 0.637 | (2,3,4,6)   |
| HD 10647†     | F8V-F9V       | 0.551  | 6151                 | -0.18$^{+0.09}_{-0.07}$ | -5.47        | 0.956 | (2,3,4,7)   |
| α Aql         | F8V           | 0.560  | 6120                 | -0.05$^{+0.09}_{-0.07}$ | -5.84        | 0.850 | (2,4,8,9,11)|
| HD 154417     | F8.5V         | 0.580  | 6042                 | -0.14$^{+0.09}_{-0.07}$ | -4.85        | 1.224 | (2,4,9,10,11)|
| HD 16673      | F6V-F8V       | 0.524  | 6301                 | -0.12$^{+0.09}_{-0.07}$ | -5.22        | 0.941 | (2,8,4,9,11)|
| HR 4767       | F8V-G0V       | 0.557  | 6038                 | -0.23$^{+0.09}_{-0.11}$ | -5.50        | 0.647 | (2,3,4,7,11)|
| ψ1 Dra B      | F8V-G0V       | 0.510* | 6223                 | -0.15$^{+0.09}_{-0.11}$ | -5.16        | 0.649 | (2,3,4,7,12)|
| HR 5581       | F7V-F8V       | 0.530  | 6156                 | -0.06$^{+0.09}_{-0.11}$ | -5.38        | 0.577 | (2,3,4,7,11)|
| HR 7793       | F7V           | 0.510** | 6261               | -0.26$^{+0.15}_{-0.11}$ | -5.44        | 0.826 | (2,4,8,9)   |
| HD 33632      | F8V           | 0.542  | 6074                 | -0.36$^{+0.09}_{-0.11}$ | -5.19        | 0.687 | (2,3,4,7,11)|

† Star with confirmed close-in giant planet (Stassun et al. 2019a).

References:–
– for spectral types and $L_X/L_{\text{bol}}$: [1] Schmitt & Liebke (2004), [6] Suchkov et al. (2003), [7] Hinkel et al. (2017), and [9] Freund et al. (2022);
– for $T_{\text{eff}}$ and [Fe/H]_{ph}: [2] Stassun et al. (2019a,b);
– for $B - V$: [11] Boro Saikia et al. (2018), [8] Baliunas et al. (1996), [12] Fabricius et al. (2002);
– $P_{\text{Rot}}$ is estimated from $B - V$ and $S_{\text{HK}}$ relation from [3] Noyes et al. (1984), or directly obtained from [5] Shkolnik et al. (2008), [8] Baliunas et al. (1996), or [10] Donahue et al. (1996);
– τC values from [4] Gunn et al. (1998) calculated using $B - V$ values is used – with $P_{\text{Rot}}$ to obtain the Rossby number $R_0$;
– [Fe/H]_{ph} measurements converted to Anders & Grevesse (1989) from [A09] Asplund et al. (2009), [G07] Grevesse et al. (2007), and [S15] Scott et al. (2015).
the rotational pole. The emission measures we estimate from the spectral analysis (see Table 5) suggest that the active area ranges from 1 - 20% of the surface area depending on the assumed coronal plasma density,\(^3\) and is thus consistent with this scenario.

Cauley et al. (2019) study SPI in the context of several models of potential interactions, specifically testing which models can generate sufficient power to generate the observed enhancement in Ca II H&K emission of several systems, including HD 179949. They found reconnection (flaring) between the star and planet was insufficient to heat H&K in many cases; instead, a model where heating is derived from field-line stretching was found more adequate to the task. Here we employ the same models to test whether the (weaker) coronal enhancement might be reconnection heated. The two main models are heating by reconnection (Lanza 2009, 2012), and heating by field line stretching (Lanza 2013). In the reconnection case (i.e., essentially flaring between star and planet as proposed by Cuntz et al. 2000), the expected power \(P_\text{s} \) is given by \(P_\text{s} = \gamma \pi / \mu R^2 B^2 / B^4 / v_\text{rel}^2\), where \(0 \leq \gamma \leq 1\) is a constant related to the relative magnetic geometry between star and planet, \(\mu\) is the magnetic permeability of the vacuum, \(R_p\) and \(B_p\) are the planetary radius and dipolar magnetic field strength, respectively, \(B_s\) is the stellar field strength at the distance of the planet, and \(v_\text{rel}\) is the relative velocity between the orbiting planet and the rotating magnetic footpoint on the star. Here, \(B_s = B_s \text{surf} (R_p / a_p)^3\), where \(B_s \text{surf}\) is the average field at the stellar surface and \(a_p\) is the orbital semi-major axis. In the field stretching case, the stellar field enters more passively, acting mostly as a conduit of energy to the star. The primary work is done on the planetary field, with power \(P_\text{s}\) given by \(P_\text{s} = 2\pi / \mu f B^2 / R^2 / B^2 / v_\text{rel}^2\). Here, \(f B_p\) is the fractional area of the planet which the magnetic footpoint occupies. Assuming \(R_p \approx R_f = 7 \times 10^9\) cm, \(\gamma = 1, B_p \approx B_f = 4\) G, \(B_s = 0.005\) G (Cauley et al. 2019), \(v_\text{rel} = 1.5 \times 10^7\) cm/s and \(f B_p = 0.1\), we have \(P_\text{s} \approx 5 \times 10^{24}\) erg/s and \(P_s \approx 7 \times 10^{27}\) erg/s. With the SPI enhancement in the X-rays of \(\Delta L_X = 4 \times 10^{26}\) ergs/s, we see that only the field stretch model (Lanza 2013) can produce the needed power for HD 179949. Thus, while the SPI appears magnetic in origin (cf. Saar et al. 2004), its estimated power points to a non-flare-like origin.

4.5. Phased variations

From Figure 10, we note a definitive correlation with the stellar polar rotation, in agreement with Scandariato et al. (2013). Apart from this, the Lomb-Scargle periodogram also suggests likely variability with the planetary orbital period, thus confirming past results of chromospheric observations (Shkolnik et al. 2003, 2005, 2008). Lastly, it is also likely that there is variability tied to the beat frequency between the stellar polar rotation and the planetary orbital period, which was suggested by Fares et al. (2012). These results can further be improved with more observations to improve phase coverage.

Note that Shkolnik et al. (2008) suggested an on/off nature of star-planet interactions, and during 2 out of 6 observation epochs, they found a periodicity of \(\sim 7\) days. We have not detected this period in our results, but non-detection of SPI in some observations is understandable, as the field stretching effect discussed earlier should occasionally cease. This could be due to decay of the active polar region on the star, or because the connection to the star is broken by excessive “field wrapping” due to the differing \(P_{\text{orb}}\) and \(P_{\text{polar}}\). In either case, some time would be needed to reestablish a sturdy star-planet connection. In Scandariato et al. (2013), their X-ray light curves also suggest variability tied to a period of \(\sim 4\) days, apart from variability with the stellar polar rotation. Due to limitations of the data used, and due to non-detection of variability tied to the planetary orbital period, they did not conclude the presence of SPI. However, we believe our result of possible variability tied to the beat period of the stellar polar rotation and the planetary orbital period agrees with their measured period of \(\sim 4\) days, and could be indicative of the presence of SPI. Alternatively, the 4 day signal could be due to two active longitudes on the star at a latitude a bit removed from the equator.

5. SUMMARY
We have carried out a comprehensive spectral and temporal analysis of the Chandra observations of HD 179949.

We first carry out spectral fits (Section 3.1) using two-temperature APEC spectral models, for which we also estimate abundances by carefully analyzing and minimizing structures in residuals using a novel technique. The resulting abundances demonstrate that HD 179949 has a larger FIP bias than is expected for stars of similar types, similar to other F stars with close-in planets. The estimated coronal Fe/H ≪ 1. While O,N/Fe ≳ 1, Ne/Fe ≪ 1.

We find that light curves in all passbands are characterized by significant variability over time scales 10^2 – 10^4 s. No flares are detected, suggesting a low-activity but dynamic environment. Hardness ratio variations are also seen, albeit at a lower significance than that of intensity, suggesting that spectral variations also occur in the corona at time scales ≲ 10^4 s.

Period analysis using the Lomb-Scargle periodogram method yields several congruous peaks that exceed 2σ significance. We find power peaks at the stellar polar rotational period (∼ 10 d), the planetary orbital period (∼ 3 d), and the beat period between the two (∼ 4.5 d). The presence of peaks at the planetary orbital time period are in agreement with Shkolnik et al. (2003, 2005, 2008). The presence of peaks at the beat frequency between the planetary orbital time period and the stellar polar rotation period suggest agreement with Fares et al. (2012) and both together are indicative of SPI, with a magnetic connection between the planetary magnetic field and the stellar magnetic equatorial field. Mechanisms of field stretching given by Lanza (2013) are explored as the possible mechanism leading to the observed SPI, as it would not always lead to chromospheric variability, thus agreeing with the on/off nature noted in Shkolnik et al. (2008), and the lack of detection of SPI in Scandariato et al. (2013). We also observe variability tied to the stellar polar rotation in agreement with Scandariato et al. (2013). The calculated area fractions provide further support for the field stretching phenomenon expected to be powering the SPI.

We also note that while HD 179949 is quite similar to other dF stars for most parameters, its photospheric metallicity [Fe/H] is higher than most of them, consistent with it having a planetary system (e.g., Fischer & Valenti 2005).

Further, the calculated F_{bias} from the coronal abundances of different metals indicates that HD 179949 shows an inverse FIP effect instead of the solar-like FIP effect expected according to Wood et al. (2018). However, they also identify stars like τ Boo A and HD 189733A which showcase an inverse FIP effect, and attribute it to the influence of their close-in Jupiter mass planet via SPI. Thus, our measured abundances provide credence to the possibility of SPI in the HD 179949 system as well.

More data are necessary to improve phase coverage and the significance of the results found here. Given the large amounts of small-scale variability in the color–color ratios, it is important to cover both soft and hard passbands using a monitoring X-ray telescope. Augmenting this with Ca II observations would be useful to test the possible physical phenomena discussed in this paper.

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Facilities: Chandra, XMM-Newton, Swift

Software: CIAO (Fruscione et al. 2006), PINTofALE (Kaspy & Drake 2000), astropy (Astropy Collaboration et al. 2022), VizieR (Ochsenbein et al. 2000)
APPENDIX

A. ADDITIONAL PLOTS

Count rate light curves for all the passbands for all ObsIDs, hardness ratio light curves for all relevant passband pairs for all ObsIDs, cumulative sum residuals plots for both models $2m$ and $2v$ for all ObsIDs are available on Zenodo (Acharya et al. 2022)$^4$. 

$^4$ 10.5281/zenodo.7220014
Figure 12. Similar to Figure 3, here we show the Cumulative sum plots of best fit models for all datasets from ObsID 5427 (top) to 6122 (bottom) as compared to the 5-95th percentile range of simulated residuals (Left). And the equivalent histogram of possible excess area values from the simulated datasets in green, the excess area of the real dataset on being fitted with the respective best fit model in red, and the 95th percentile of excess areas as a dashed grey line (Right). Note that all accepted models have an excess area less than the 95th percentile limit.
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