IS IT JUSTIFIED TO ASSUME THAT "EVERYWHERE IN THE SUN'S PHOTOSPHERE-CORONA DOMAIN THE ELECTRIC CONDUCTIVITY IS HIGH"?; OR, WHAT DRIVES THE SOLAR UPPER ATMOSPHERE?

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ABSTRACT

For many years, solar scientists have recognized the extreme complexity of the upper solar atmosphere. However, in order to construct a valid theoretical model they have formulated a set of simplified assumptions governing the makeup of the Sun's upper atmosphere. In previous papers (Feldman 1983; Feldman 1993) a number of assumptions used to construct the models were shown not to be valid. In this Paper I bring evidence to question the validity of the last and most important of the assumptions, the assertion that "Everywhere in the photosphere-corona domain the electric conductivity is HIGH." The consequences of the finding are briefly discussed. A laboratory measurement of the electrical conductivity versus temperature in a gas with photospheric composition would resolve this important issue and seem to be feasible.

Subject headings: MHD — Sun: atmosphere — Sun: corona

1. INTRODUCTION

The modern era of solar coronal research started some 50 years ago with the identification of a number of peculiar lines recorded in the visible region of the spectrum during solar eclipses. Edlén (1942) showed that the visible lines originate from forbidden transitions within ground terms of highly ionized atoms (i.e., Fe x, Fe xi, Ca xii, and Ca xiii). The presence of such highly ionized species established that the coronal temperature is in the 10⁶ K range. To this day, the explanation for the mechanism which fuels a corona, some 200 times hotter than the photosphere, has eluded us.

Although it is realized by many that the solar atmosphere is very complex, a number of simplified assumptions have been formulated to enable the development of a realistic theoretical model for the solar atmosphere. Some of the most important of these assumptions are given below.

ASSUMPTION I. The upper solar atmosphere can be approximated by "plane parallel formations."

ASSUMPTION II. The temperature \(T_e\) is a monotonically increasing function of the radius vector \(R\).

ASSUMPTION III. The elemental abundances in the upper solar atmosphere and in the photosphere are identical.

ASSUMPTION IV. The plasma is in steady state coronal equilibrium.

ASSUMPTION V. The plasma is in constant pressure \((T_e \times N_e) = \text{constant}\).

ASSUMPTION VI. Everywhere in the photosphere-corona domain the electric conductivity is HIGH.

In spite of the fact that solar modelers recognized that a number of the above assumptions are too simplistic, to this day some and, on occasion, all of the above assumptions are being used by solar and stellar physicists alike in the construction of upper atmosphere models for the Sun and for solar-type stars. Ironically, in all of the six assumptions listed above, the truth is often closer to the opposite of what has been assumed. Although this paper is devoted to Assumption VI, for the sake of completeness and for the benefit of readers not familiar with the details of the solar upper atmosphere, arguments against the first five assumptions will be briefly mentioned in the Introduction.

It is only necessary to look at pictures of the solar atmosphere formed at the 10⁵ K (Fig. 1) and the 10⁶ K (Fig. 2) ranges to understand that the first assumption is not valid not only in a full Sun framework but also in the context of fairly small solar regions. The first assumption should have been written:

ASSUMPTION I. The upper solar atmosphere CANNOT be approximated by "plane parallel formations."

The second assumption claims that a continuity exists between the cold chromosphere and the hot corona in which the temperature increases in a monotonic fashion. An important ingredient of the model is the thin transition region (Fig. 3). As a result of experiments on Skylab it was shown (Huber et al. 1974; Feldman 1983) that the radiation attributed to the so-called transition region originates not from a very thin layer situated between the chromosphere and corona but mostly from an entity composed of numerous small structures. At an average height of \(\approx 3000\) km above the white-light limb, a region where they are most abundant, the structures occupy only \(\approx 1\%\) of the volume (Feldman, Doschek, & Mariska 1979). Feldman (1983) suggested that the so-called transition region should be renamed "unresolved fine structures." By studying monochromatic solar images taken by the Skylab SO82a (e.g., An Atlas of Extreme Ultraviolet Spectroheliograms from 170 to 625 Å; Feldman, Purcell, & Dohne 1987), it can be seen that cold coronal regions often extend to heights much larger than nearby hotter regions. The only thing that can be said regarding the second assumption is that, in general, starting at \(\approx 3000\) km above the white-light limb the ratio of hot to cold plasma increases as one moves higher and higher in the solar upper atmosphere. Thus, the second assumption should be stated as follows:

ASSUMPTION II. The temperature \(T_e\) is NOT a monotonically increasing function of the radius vector \(R\). In general, at a radius greater than \(\approx 3000\) km above the white light limb the ratio of hot to cold plasma increases with height.
The research of elemental abundances in the upper solar atmosphere is fairly extensive. At some point it was assumed that because of large temperature gradients a mass-dependent separation of elements should take place (i.e., Delache 1967). However, since no such effect was ever discovered and because of uncertainties in the atomic data, it was assumed that elemental abundances in the upper solar atmosphere and in the photosphere are identical (see also § 2.3). Elemental abundances in the solar wind (SW) and in solar energetic particles (SEP) were known to differ from those in the photosphere (Breneman & Stone 1985; Bochsler, Geiss, & Kunz 1986; Gloeckler & Geiss 1989; Coplan et al. 1990; and Stone 1989) in a manner shown schematically in Figure 4. In recent years, it became apparent that coronal elemental abundances also vary relative to those in the photosphere. Specifically, Meyer (1985) has shown that, on the average, when compared with the photosphere, upper solar atmosphere elements with first ionization potential (FIP) < 10 eV are enhanced by factors of 3–4. A more detailed study revealed that, in general, the abundance of low-FIP elements can be either photospheric, enhanced by a factor of 20 relative to the photosphere or can acquire any value between these two extremes (Widing & Feldman 1989).
The amount of variation depends on properties of the emitting region. For a review on the properties of elemental abundances in the upper solar atmosphere see Feldman (1992a). Thus, the third assumption should be rephrased as follows:

ASSUMPTION III. The elemental abundances in the upper solar atmosphere and in the photosphere are NOT necessarily identical. They may be the same or vary by any factor between 1 and 20.

It is assumed by most that the upper solar atmosphere is in steady state equilibrium, implying that plasma conditions do not change in time which are of the same order as the ionization and recombination times of the ions in question. Based on this assumption, many atomic systems have been analyzed and understood resulting in a significant body of knowledge; however, some atomic systems could not be understood in the context of a steady state equilibrium. Recently, it was shown that some of the most problematic atomic systems are easily explained in the context of transiently ionizing plasmas. When compared with line intensity calculations, which are based on the assumption of steady state coronal equilibrium, the measured absolute intensity of He II lines appear to be an order of magnitude too large, and the relative intensities within the Lyman series are incorrect. Jordan (1975) showed that the discrepancy between calculations and observations can be resolved by assuming that He II ions diffuse across the thin transition region into a hotter domain, a region where they become excited. Since the thin transition region idea was shown to be invalid, a different explanation needs to be found. Table 1 shows results from data acquired during a flare. By raising the temperature of the emitting plasma from \( T_e \approx 8 \times 10^6 \) K, a temperature at which the He II lines are emitted in steady state coronal equilibrium, to \( T_e \approx 2 \times 10^7 \) K, calculations of the relative and the absolute intensities match the experimental values (Laming & Feldman 1992). However, at such temperatures, He II ions will rapidly ionize. Similar problems with Fe R (\( T_e \approx 1 \times 10^6 \) K) density-sensitive line ratios (Feldman 1992b), the Fe XV (\( T_e \approx 2 \times 10^6 \) K) \((3s^21S_0-3s3p^2P_J)/(3s^21S_0-3s3p^1P_J)\) intensity ratio (Feldman et al. 1992) and the Fe xxv/Fe xxiv intensity ratio (\( T_e \approx 2 \times 10^7 \) K) (Koshelev & Kononov 1981) can be resolved by assuming a transiently ionizing plasma. It seems apparent that the fourth assumption should be rewritten as follows:

ASSUMPTION IV. The upper solar atmosphere plasma, often, is NOT in steady state coronal equilibrium, but in a state of transient ionization.

In order to check the assumption of constant pressure \( (N_e \times T_e = \text{constant}) \) in the upper solar atmosphere, it is required to locate plasma structures that extend over a wide temperature range. The unresolved fine structures fulfill the above requirements. The solar spectrum in the 1175–1930 Å wavelength range was thoroughly observed by the Skylab normal incidence spectrometer SO82b (Bartoe et al. 1977). Spectra obtained by this instrument contain a number of intense lines which are emitted by \( 5 \times 10^4 \leq T_e \leq 2.5 \times 10^5 \) K plasmas and are good electron density indicators for solar plasma (Fig. 5). Since the lines originate from low-lying levels, their properties are not affected by the assumed excitation temperature. However, since they span a factor of 5 in temperature, they can be used to check if, on the average, they are emitted by plasmas in constant pressure \( (N_e \times T_e = \text{constant}) \), as may be the case for a loop in a steady state coronal equilibrium, or from plasmas with constant density, as may be the case in plasmas undergoing temperature changes which are much faster than the changes in volume.

Feldman & Laming (1993) investigated electron densities associated with quiet, active, and flare regions in the \( 5 \times 10^4 \leq T_e \leq 2.5 \times 10^5 \) K range using the above-mentioned lines. They have concluded that these particular \( 5 \times 10^4 \leq T_e \leq 2.5 \times 10^5 \) K plasma regions can be described better as

<table>
<thead>
<tr>
<th>Flare</th>
<th>Transition</th>
<th>Data</th>
<th>Fit</th>
</tr>
</thead>
<tbody>
<tr>
<td>Dec 22</td>
<td>1s–4p</td>
<td>2.55 \times 10^{13}</td>
<td>2.50 \times 10^{13}</td>
</tr>
<tr>
<td></td>
<td>1s–5p</td>
<td>1.64 \times 10^{13}</td>
<td>1.72 \times 10^{13}</td>
</tr>
<tr>
<td></td>
<td>1s–6p</td>
<td>1.44 \times 10^{13}</td>
<td>1.25 \times 10^{13}</td>
</tr>
<tr>
<td></td>
<td>1s–7p</td>
<td>9.84 \times 10^{14}</td>
<td>1.01 \times 10^{15}</td>
</tr>
<tr>
<td></td>
<td>1s–8p</td>
<td>7.46 \times 10^{14}</td>
<td>8.57 \times 10^{14}</td>
</tr>
<tr>
<td>Jan 21</td>
<td>1s–4p</td>
<td>1.23 \times 10^{15}</td>
<td>1.11 \times 10^{15}</td>
</tr>
<tr>
<td></td>
<td>1s–5p</td>
<td>6.44 \times 10^{14}</td>
<td>7.01 \times 10^{14}</td>
</tr>
<tr>
<td></td>
<td>1s–7p</td>
<td>3.62 \times 10^{14}</td>
<td>3.92 \times 10^{14}</td>
</tr>
<tr>
<td></td>
<td>1s–8p</td>
<td>3.27 \times 10^{14}</td>
<td>3.24 \times 10^{14}</td>
</tr>
<tr>
<td></td>
<td>1s–9p</td>
<td>2.90 \times 10^{14}</td>
<td>2.69 \times 10^{14}</td>
</tr>
</tbody>
</table>

Note.—The data and the fits are in units of photons \( s^{-1} \text{ cm}^{-2} \) emitted into 4π solid angle.

![Fig. 5.—Intensity ratios of spin-forbidden/allowed lines formed in the \( 5 \times 10^4 \leq T_e \leq 2.5 \times 10^5 \) K temperature range.](image-url)
TABLE 2

<table>
<thead>
<tr>
<th>Ion</th>
<th>Line Ratio</th>
<th>( T_e (K) )</th>
<th>Height Above the Solar Limb</th>
</tr>
</thead>
<tbody>
<tr>
<td></td>
<td></td>
<td></td>
<td>-12°</td>
</tr>
<tr>
<td>Si ii</td>
<td>1301/1296</td>
<td>( 5 \times 10^4 )</td>
<td>11.4</td>
</tr>
<tr>
<td>Si ii</td>
<td>1301/1303</td>
<td>( 5 \times 10^4 )</td>
<td>11.2</td>
</tr>
<tr>
<td>C ii / C iv</td>
<td>1909/1403</td>
<td>( 7 \times 10^4 )</td>
<td>…</td>
</tr>
<tr>
<td>S iv</td>
<td>1416/1406</td>
<td>( 9 \times 10^4 )</td>
<td>11.0</td>
</tr>
<tr>
<td>O iv / C iv</td>
<td>1666/1551</td>
<td>( \sim 1 \times 10^5 )</td>
<td>11.1</td>
</tr>
<tr>
<td>O vii / N v</td>
<td>1401/1240</td>
<td>( \sim 1.6 \times 10^5 )</td>
<td>11.3</td>
</tr>
<tr>
<td>O vii / N v</td>
<td>1218/1240</td>
<td>( \sim 2.5 \times 10^5 )</td>
<td>11.2</td>
</tr>
</tbody>
</table>

* From Table 2 in Feldman & Laming 1993.

Assumption V. The plasma is NOT always in constant pressure \((T_e N_e = \text{constant})\). Most likely, while undergoing a heating process, individual volumes maintain a nearly constant density.

A detailed discussion of the arguments against Assumptions IV and V were given recently in a talk presented at the Nobel Symposium 85 (Feldman 1993, hereafter Paper I). This paper can be considered to be Paper II in the series in which Assumption VI, which is most difficult to evaluate, is being discussed. The rest of the paper will be devoted to the task of showing possible flaws in Assumption VI and their implications.

2. HOW VALID IS THE ASSUMPTION THAT "EVERYWHERE IN THE PHOTOSPHERE-CORONA DOMAIN THE ELECTRIC CONDUCTIVITY IS HIGH"?

Perhaps the principal hypothesis steering coronal heating research to solutions involving magnetohydrodynamic processes is the notion that the solar atmosphere is very conductive, from the lower layer of the photosphere through the corona. Parker (1979) articulates this idea as follows:

But it appears that everywhere in the universe there are enough x-rays, ultraviolet radiation, cosmic rays, and thermal excitation to guarantee that at least a few electrons everywhere will be dislodged from the local atoms at any time. Certainly this is the rule in the Sun, where even in the cool photosphere the carbon and the alkali metals are at least partially ionized. Elsewhere in the Sun the gases are fully ionized and saturated with free electrons, … The free electrons quickly neutralize any separated charges. The characteristic time to carry out the neutralization is the plasma period \((n_{e} / N_{e} e^2)^{1/2} \text{ s}\) where \(n_{e}\) is the electron mass, \(e\) the charge, and \(N_{e}\) the number of free electrons per cm\(^3\). Numerically this is \(10^{-4} / N^{1/2} \text{ s}\). Even in intergalactic space where \(N_{e}\) may be \(10^{-10} - 10^{-8} \text{ cm}^3\), the neutralization time is only \(1 - 10 \text{ s}\). Elsewhere it is much less. It follows, then, that any large-scale electrostatic field is quickly and effectively neutralized by the universal free electrons. The general occurrence of electric field is restricted to microscopic scales, inside atoms or within a Debye radius of an ion. The large scale electric field in the frame of the gas containing the free thermal electrons is essentially zero.

Obviously all that Parker has said is correct provided that the information supplied by solar modelers stating that “the minimum temperature of the solar atmosphere is high enough and the radiation field is sufficiently intense to keep a sizable fraction of the atoms in every part of the Sun ionized” is reliable. Deductions like the one voiced by Parker (1979) resulted in channeling efforts to find the source of coronal heating away from processes resulting from regions containing large charge separations, i.e., large-scale potential differences and high electrical currents, and into magnetohydrodynamics.

A careful evaluation of new and old experimental results suggests that there are good reasons to believe that in some parts of the solar atmosphere the ratio of electrons to neutral atoms is three or more orders of magnitude smaller than previously believed. In the context of the above sentence, Assumption VI stating that Everywhere in the photosphere-corona domain the electric conductivity is HIGH, may become questionable.

2.1. Elemental Abundances and the Ratio of Free Electrons to the Number of Neutral H Atoms \(N_e / N_H\) in Cold Solar Regions

A study of the solar photosphere composition reveals that the relative abundance of the total number of particles with a FIP less than a specific value is a very sensitive function of the FIP-value. Table 3 lists all the elements in the periodic table for which their photospheric elemental abundance relative to H exceeds \(1 \times 10^{-7}\). The photospheric elemental abundances used are from the compilation by Anders & Grevesse (1989). Table 3 is arranged according to the FIP starting at He (FIP = 24.6 eV) and ending with K (FIP = 4.3 eV). Notice that the most abundant elements in the photosphere have a FIP > 11 eV. The total number of atoms relative to H + He having FIPs < 10 eV is only \(1.1 \times 10^{-4}\), for those having FIPs < 7 eV the number is only \(7.2 \times 10^{-6}\), for those with FIPs < 6 eV the number is \(2.0 \times 10^{-6}\) and for those with FIPs < 5 eV the number is only \(1.3 \times 10^{-7}\) (see Fig. 6). In the interval between the FIP of C (11.5 eV) to a FIP of smaller than that of Na (5.2 eV) the abundance ratio relative to H is reduced by 3.5 orders of magnitude. The conductivity of a mildly ionized plasma where the density of free electrons to the number density of neutral H + He atoms \(N_{e} / N_{H+He}\) changes from \(4.5 \times 10^{-4}\) to 1.3 \(\times 10^{-7}\) will be significantly modified. It is important to note that the only elements in the periodic table with FIP < 4.3 eV are Rb (FIP = 4.2 eV) and Cs (FIP = 3.9 eV) and their abundances relative to H are \(2.5 \times 10^{-10}\) and \(1.3 \times 10^{-11}\) respectively.
TABLE 3

LIST OF THE MOST ABUNDANT ELEMENTS IN THE SOLAR PHOTOSPHERE

<table>
<thead>
<tr>
<th>Element</th>
<th>Z</th>
<th>FIP (eV)</th>
<th>Photospheric Abundance</th>
<th>&quot;Coronal&quot; Enrichment</th>
</tr>
</thead>
<tbody>
<tr>
<td>He</td>
<td>2</td>
<td>24.6</td>
<td>$1.0 \times 10^{-1}$</td>
<td></td>
</tr>
<tr>
<td>Ne</td>
<td>10</td>
<td>21.6</td>
<td>$1.2 \times 10^{-4}$</td>
<td>1</td>
</tr>
<tr>
<td>Ar</td>
<td>18</td>
<td>15.8</td>
<td>$3.6 \times 10^{-6}$</td>
<td>1</td>
</tr>
<tr>
<td>N</td>
<td>7</td>
<td>14.5</td>
<td>$1.1 \times 10^{-4}$</td>
<td>1</td>
</tr>
<tr>
<td>H</td>
<td>1</td>
<td>13.6</td>
<td></td>
<td>1</td>
</tr>
<tr>
<td>O</td>
<td>8</td>
<td>13.6</td>
<td>$8.5 \times 10^{-4}$</td>
<td>1</td>
</tr>
<tr>
<td>Cl</td>
<td>17</td>
<td>13.0</td>
<td>$1.9 \times 10^{-7}$</td>
<td>1</td>
</tr>
<tr>
<td>C</td>
<td>6</td>
<td>11.3</td>
<td>$3.6 \times 10^{-4}$</td>
<td>1</td>
</tr>
<tr>
<td>P</td>
<td>15</td>
<td>10.5</td>
<td>$3.7 \times 10^{-7}$</td>
<td>1</td>
</tr>
<tr>
<td>S</td>
<td>16</td>
<td>10.3</td>
<td>$1.9 \times 10^{-5}$</td>
<td>1.5</td>
</tr>
<tr>
<td>Si</td>
<td>14</td>
<td>8.1</td>
<td>$3.6 \times 10^{-5}$</td>
<td>3</td>
</tr>
<tr>
<td>Fe</td>
<td>26</td>
<td>7.9</td>
<td>$3.2 \times 10^{-5}$</td>
<td>3</td>
</tr>
<tr>
<td>Mg</td>
<td>12</td>
<td>7.6</td>
<td>$3.8 \times 10^{-5}$</td>
<td>3</td>
</tr>
<tr>
<td>Ni</td>
<td>28</td>
<td>7.6</td>
<td>$1.8 \times 10^{-6}$</td>
<td>3</td>
</tr>
<tr>
<td>Mn</td>
<td>25</td>
<td>7.4</td>
<td>$3.4 \times 10^{-7}$</td>
<td>5</td>
</tr>
<tr>
<td>Cr</td>
<td>24</td>
<td>6.8</td>
<td>$4.8 \times 10^{-7}$</td>
<td>5</td>
</tr>
<tr>
<td>Ca</td>
<td>20</td>
<td>6.1</td>
<td>$2.2 \times 10^{-6}$</td>
<td>5</td>
</tr>
<tr>
<td>Al</td>
<td>13</td>
<td>6.0</td>
<td>$3.0 \times 10^{-6}$</td>
<td>6</td>
</tr>
<tr>
<td>Na</td>
<td>11</td>
<td>5.2</td>
<td>$2.1 \times 10^{-6}$</td>
<td>6</td>
</tr>
<tr>
<td>K</td>
<td>19</td>
<td>4.3</td>
<td>$1.4 \times 10^{-7}$</td>
<td>9</td>
</tr>
</tbody>
</table>

Note.—Elements are arranged according to their (FIP).


* Enrichment values are estimates for the average Sun. The values for Cr, Al, and K are less reliable than the other values.

2.2. Measurements and Models of Cold Regions in the Lower Solar Atmosphere

The traditional one-component model of the lower solar atmosphere (i.e., VAL by Vernazza, Avrett, & Loeser 1981) attempted to determine, among other things, the value of $N_e/N_H$. The models utilize measured intensities from optically thin and thick lines, continuum intensity measurements in the 1600 Å region of the spectrum and the photospheric elemental abundances. According to such models, the minimum temperatures in the lower solar atmosphere are usually in excess of 4000 K. The models also predict that at such temperatures Mg, Si, and Fe will be substantially ionized while the higher FIP elements, of which some are more abundant, are mostly neutral. Since the abundance of these three elements relative to H is $\approx 10^{-4}$ (see Table 3 and Fig. 6) the relative number of free electrons is $N_e/H = He \approx 10^{-4}$.

Grevesse & Sauval (1991) summarized models of the lower solar atmosphere which incorporate infrared observations of molecular lines. According to them, Ayres (1981), Ayres, Tasterman, & Brault (1986), and Ayres (1990) have convincingly shown that the empirical one-component models of the solar outer layers (Maltby et al. [1986]; VALC) and Holwege & Müller (1974; HM) fail to explain the very optically thick lines, i.e., Lyα, Ca II, H and K as well as the infrared CO lines at the center of the disk and their center-to-limb variations. In order to resolve the discrepancy, Ayres (1981) proposed a thermal bifurcation model for the solar atmosphere composed of two thermally distinct zones of hot and cool matter which coexist at the same altitude. Ayres et al. (1986) found that the filling factor of the surface coverage of the hot regions should be less than 10% at the level of formation of the CO lines, where the cool matter which contains the CO occupies more than 90% of the solar surface. Ayres et al. (1986) showed that his two-component model (COOLC and FLUXT; Fig. 7) agrees with observations of the optically thin lines as well as with the CO observations. Grevesse & Sauval (1991) have constructed a new model (GSCO; Fig. 7) based on high-resolution solar spectra obtained by a Fourier transform spectrometer experiment (ATMOS) on board the Space Shuttle which provided a set of improved infrared observations of molecular lines. Notice that the GSCO model predicts temperatures significantly

![Fig. 6.](image_url)

![Fig. 7.](image_url)
lower than those predicted by Ayres et al. (1986; COOLC). A theoretical non-LTE model in radiative equilibrium by Anderson (1989; AND) looks similar in shape to the COOLC and GSCO semiempirical models; however, it predicts still lower temperatures ($T_e \approx 2700$ K) in the cooler parts of the lower solar atmosphere. It is conceivable that in some distinct regions of the solar atmosphere the actual temperatures can be even lower than those predicted by Anderson's model.

Solar atmosphere regions where $T_e \lesssim 2700$ K will contain a significantly lower fraction of electrons than regions predicted by earlier models like the VAL. In fact, if there are cold enough regions where the majority of ions with FIP $> 5$ eV are neutral, the ratio of electrons to H atoms can be only $N_e/N_{H+H_2} \lesssim 1 \times 10^{-7}$. The electric conductivity of such a region will certainly be significantly reduced from those predicted by the earlier models. If it will be discovered that in some solar regions even K is mostly neutral, the fraction of free electrons relative to neutral atoms may be immeasurably small. As was shown by Jordan et al. (1978) and Bartoo et al. (1979), the H$_2$ molecule is rather abundant in sunspots. The dissociation energy of H$_2$ (4.48 eV) is similar to the FIP of K.

2.3. Elemental Abundance Variations in Low-FIP Atoms

Mg, Si, and Fe occupy a special place among elements abundant in the solar atmosphere. They are the most abundant among the low-FIP elements with almost equal abundances and FIPs (see Table 3). The only real difference among them are their atomic weights. While the atomic weights of Mg and Si are 24 and 28, the atomic weight of Fe is 56, about 2.3 times larger than the atomic weight of Mg and 2 times the atomic weight of Si. As far as we can determine, for any region observed, the upper solar atmosphere enrichment of Mg, Si, and Fe seems to be the same.

McKenzie & Feldman (1992) measured the Fe XVII/Mg XI ratios in a large number of flares and displayed them versus the temperature-sensitive Fe XVI/Fe XVII line ratios (see Fig. 8a). An Fe/Mg abundance ratio of 0.8 was determined from the line ratios. The photospheric ratio given by Anders & Grevesse (1989) is 0.84. Because of the poor statistics of the line ratios, McKenzie & Feldman (1992) have concluded that, if any, the actual Fe XVII/Mg XI abundance variations do not exceed a factor of 1.5. For comparison purposes the ratios of the low-FIP Fe XVII to the high-FIP Ne X line obtained from the same flares are shown in Figure 8b.

The 417.25 Å (3s$^2$ 1S$_0$–3s3p$^3$P$^o$) Fe XV line and the 520.66 Å (2s$^2$ 2S$_{12}$–2p$^2$ 2P$^{o1}$) Si XIX are emitted at almost the same temperature as seen by their contribution functions in Figure 9. The two lines are quite intense and can be seen at high heights above active regions. They are also very prominent in the hot coronal clouds. A visual inspection was conducted on images published in the SO82a Atlas (Feldman et al. 1987) and later on the entire SO82a data base to search for coronal clouds in which the Fe XV/Si XIX line ratios are different from the expected values. The spatial resolution of the inspected images was better than 2000 km. During the entire visual search, which included a large number of active regions, some of which reached heights and widths of well over $1 \times 10^4$ km, no case was found where there was a significant variation in the Fe XV/Si XIX ratio either between different clouds or within a single cloud. It is estimated that the eye could have easily detected variations as large as a factor of 1.5. It is safe to conclude from the constancy of either the Fe XV/Mg XI or Fe XV/Si XIX ratios discussed above that the effect of atomic mass variation on elemental abundances is very small or nonexistent. Thus, abundance variations result primarily from variations in FIPs.

A dilemma is posed by the Ca abundance in the upper solar atmosphere. Mg, Ca, and Fe are defined as low-FIP elements. The photospheric abundance ratio of Ca to Fe is $\sim 0.071$ and that of Ca to Mg is $\sim 0.059$ (Anders & Grevesse 1989). Measurements from solar flares with temperatures of $\approx 2 \times 10^7$ K (i.e., Doschek, Feldman, & Seely 1985; Phillips & Feldman 1991) indicate Ca/Fe ratios of 0.10 and 0.16 which are 1.4 and 2.2 times larger than the photospheric values. Measurements from plasmas in the temperature range $T_e \sim 5 \times 10^7$ K emitted either by an impulsive flare (Feldman & Widing 1990) or from a polar plume (Widing & Feldman 1992) also indicate Ca/Mg ratios of 0.10 and 0.11 which are 1.7 and 1.9 times larger, respectively, than the expected photospheric value.

The photospheric ratio of Na/Mg is 0.054. Feldman & Widing (1990) have measured a ratio of 0.13 in an impulsive flare and Widing & Feldman (1992) have measured a ratio of 0.11 and 0.077 at different heights of a polar plume. The measured values are 2.4, 2.0, and 1.4 times larger than expected in the photosphere.

In order to determine the abundance of K in the solar upper atmosphere Doschek et al. (1985) measured the K XVIII/Ca XIX line intensity ratio in a spectrum obtained by the NRL P78-I Softflex experiment. The spectrally used for the measurement was obtained by summing a large number of spectral scans in order to improve the signal-to-noise ratio and detect weak lines that might not be noticed in a single scan. In order to remove possible spectral variations with time from the summed results, spectra were selected from scans near the time of peak flare emission, when plasma conditions are changing rather slowly, for selected time intervals during the decay phase, and for flares with relatively long decay times. Since the K XVIII and Ca XIX lines arise from similar He-like transitions with similar energies, the measured ratio has a negligible temperature dependency. According to Doschek et al. (1985) the K/Ca abundance ratio is 0.10. The K/Ca flare data is a factor of 1.7 higher than the photospheric ratio K/Ca = 0.058 given by Anders & Grevesse (1989).

McKenzie & Feldman (1993) have measured abundances of Ca and Cr using soft X-ray lines ($\lambda < 25$ Å) which are emitted by active region plasmas. They have determined that the coronal abundance ratio of Cr/O is probably always enhanced relative to the photosphere by at least a factor of 3. The Ca abundances when compared with the Fe abundance were enhanced on the average by a factor of 2.

If the elemental abundance measurements of Cr, Ca, Al, Na, and K are correct, and there is no reason to suspect that they are not, it will have to be concluded that the enrichment process of the upper solar atmosphere is not a single-step process but a continuous one. Thus, we will have to conclude that the level of enrichment of low-FIP elements in the upper solar atmosphere depends on the FIP's magnitude. The lower the FIP, the larger is the level of enrichment, which increases with declining FIP. Figure 10 shows the average level of enrichment relative to the Mg, Si, and Fe for low-FIP elements. The scatter in the data may be due not only to statistical uncertainties in the measurements, temperature variations, or uncertainties in the atomic data, but also to actual changes in the environment.
Fig. 8—(a) The photon flux ratio of the Fe xvm $2p^4 2P_{3/2}-2P_{1/2} (^1D)$ 3d $2D_{3/2}-2P_{3/2}$ line blend to the Mg xi $1s^2 1S-1s^2 2p^1 P$ line plotted as a function of the ratio of the Fe xvm to Fe xvim flux. The abscissa is nonlinear. The temperatures on the top scale are those for which an isothermal plasma will have the corresponding Fe xvm/Fe xvim photon flux ratios. The curve drawn through the data is the expected value when the theoretical abundance ratio from Anders & Grevesse (1989) is used. Except for the low-temperature end, the theoretical ratio is independent of temperature. (b) The photon flux ratio of the Fe xvm line to the Ne ix $1s^2 1S-1s^2 2p^5 S$ forbidden line plotted as a function of the ratio Fe xvm to Fe xvim flux (McKenzie & Feldman 1992).
SOLAR UPPER ATMOSPHERE

2.3.1. A Model Describing the Low FIP Ion Enrichment Process in the Upper Solar Atmosphere

The fact that enrichment of the upper solar atmosphere is a function of the FIP implies that it operates in interfaces between a colder and a hotter region where higher FIP ions are neutral on the colder side of the transition and ionized in the hotter, while low FIP ions are continuously being generated on the colder side and being transferred into the hotter side. Thus, since we see a difference in abundance between Mg, Si, and Fe (FIP \( \approx 8 \) eV) and the higher FIP ions (FIP \( > 11 \) eV), it is expected that there are regions where Mg, Si, and Fe atoms are becoming ionized but where atoms belonging to higher FIP elements are staying neutral. By the same token, since we see a difference between Na (FIP = 5.2 eV) and Ca (FIP = 6.1 eV) on the one end and the higher “low” FIP Mg, Si, and Fe (FIP \( \approx 8.0 \) eV) on the other, we expect to find regions where Mg, Si, and Fe atoms are neutral while Ca, Na, and K atoms are being ionized. The difference in the K/Ca abundance ratio between the photosphere and the upper solar atmosphere suggests that we can go one step further and postulate the existence of cold regions in which only K (FIP = 4.3 eV) atoms are ionized but Ca (FIP = 6.1 eV), and possibly, Na (FIP = 5.2 eV) atoms are mostly neutral. These may be the regions from which the cold CO molecular radiation is being emitted (see § 2.1). The electron to neutral H density ratio in regions where only the very low FIP \( < 5 \) eV are becoming ionized will be \( N_e/N_{H+He} \leq 1 \times 10^{-7} \). If it will be found that the solar atmosphere contains regions that even K is not ionized, the \( N_e/N_{H+He} \) will be negligibly small.

The enrichment model assumes the presence of sufficiently strong electric fields which act on the interface between the hotter and colder regions and transfer ions into the hotter domains. Under this model, the rate of the removal of ions from the colder into the hotter region must be larger than the rate of ion creation in the colder regions. Under such conditions, all ions created in the cold region, in places where the electromagnetic field is effective, are transferred into the hot region, and no mass-dependent enrichment by low-FIP ions is expected. K, Na, Al, and Ca, which are being ionized at a lower temperature than Mg, Si, and Fe, are expected to be enriched in the upper solar atmosphere, when compared to the photosphere, by a larger ratio than the higher FIP ions Mg, Si, and Fe.

2.4. Summary

In § 2.1 it was shown that the total fraction of particles with a FIP smaller than a certain value is a steep function of the FIP (Fig. 6). By lowering the FIP from 8.1 to 5.1 eV, the total fraction of particles decreases by three orders of magnitude. Thus, in plasmas where elements with FIP > 5 are mostly not ionized, the ratio of electrons to neutral H + He atoms will be \( N_e/N_{H+He} \leq 1 \times 10^{-7} \).

Section 2.2 summarized observations and calculations partly semiempirical and partly theoretical, indicating that the temperature in large regions of the lower solar atmosphere may be significantly lower than previously believed \( (T_e \approx 2700 \) K). It is quite conceivable that in some regions the temperature can be perhaps even lower.

Section 2.3 described measurements showing that there are good reasons to believe that the elemental abundance ratios of K/Ca and Ca/Fe differ in the upper solar atmosphere from their value in the photosphere. Such results can be explained by postulating the existence of lower solar atmosphere regions where Fe and elements with similar FIP values are not ionized, moreover, some other colder regions where Ca and similar FIP elements are also not ionized. These are indications that the quantity \( N_e/N_{H+He} \) in parts of the lower atmosphere may be low. Moreover, the entire process of enrichment of the upper solar atmosphere can be reasonably explained by the presence of electric fields between hotter and colder regions of the lower solar atmosphere, providing a place for the enrichment process to occur.
3. EXPERIMENTAL EVIDENCE TO SUGGEST THE PRESENCE OF LARGE ELECTRICAL CURRENTS IN THE UPPER SOLAR ATMOSPHERE

Direct measurements of magnetic fields in hot regions of the upper solar atmosphere are difficult and have not yet been performed. The difficulty rests with the fact that high-temperature coronal lines have relatively large Doppler widths, which obscure the shifts of the Zeeman components required for the measurements. Below, I discuss evidence to suggest that plasma is being compressed in regions of the upper solar atmosphere, a process that may be attributed to plasma columns containing strong azimuthal fields which are driven by large electric currents.

3.1. Electron Density in Unresolved Fine Structures \( T_e \approx 5 \times 10^4 \) K

As was already discussed in the Introduction, the solar spectrum in the 1175–1930 Å wavelength range contains a number of density-sensitive lines. In particular, the upper level population of the C III 1909 Å intersystem line reaches a quasi-statistical equilibrium with the ground state which does not vary with electron densities \( (1 \times 10^4 < N_e < 1 \times 10^5 \text{ cm}^{-3}) \) typical in the upper solar atmosphere (see Fig. 5). When evaluated for different heights and for regions having variations in brightness, the C III 1909 Å line intensities provide a measure of the number of particles \( N_e V \) along the line of sight for each of the emitting volumes (V) (Fig. 11a). Figure 11b shows the Si iv 1403 Å line intensities in identical height locations and regions. Being an allowed line, its intensity is proportional not only to the number of particles but also to the electron density, i.e., to \( N_e \times (N_e V) = N_e^2 V \). Thus the Si iv 1403 Å/C III 1909 Å intensity ratios are proportional to the electron densities in the regions (Fig. 11c). For details on the solar regions and their properties responsible for the C III and Si iv observations, see Feldman & Doschek (1978).

Although the amount of material in the different regions is the same within a factor of 2 (Fig. 11a), the emitted intensity in allowed lines from the different parts of the active regions change by as much as a factor of 20 (Fig. 11b). The changes are caused by the density increasing from \( N_e \approx 1 \times 10^{10} \text{ cm}^{-3} \) in the typical quiet regions to \( 2 \times 10^{11} \text{ cm}^{-3} \) in dense active regions.

3.2. Electron Density in Flaring Regions \( T_e \approx 2 \times 10^6 \) K

Gabriel & Jordan (1969) have described an elegant method for determining electron densities in the 2 x 10^6 K temperature range using the forbidden to intercombination intensity ratios in the O vii lines \( (1s^2 3S_0-1s2s^2 3S_1/1s^2 3S_0-1s2p^2 3P_j) \) near 22 Å. The sensitivity to electron densities larger than \( N_e > 1 \times 10^{10} \text{ cm}^{-3} \) (see Fig. 12) occurs primarily because the population of the 1s2s^2 3S_1 level approaches a quasi-equilibrium with the 1s^2 3S_0 ground state at those densities. Consequently, the forbidden-line intensity reflects mainly the variations in the number of O vii ions along the line of sight, much like the C III 1909 Å line does. On the other hand, under solar flare plasma conditions, the intercombination line behaves much like the allowed Si iv 1404 Å line (see § 3.1). Figure 13 presents the O viii line intensities near 22 Å before the start and during the 1980 May 9 solar flare (Doschek et al. 1981). Notice that, prior to the flare onset \( (t < 12 \text{ minutes } 56 \text{ s}) \) the ratio between the forbidden and intercombination lines is nearly 1, while immediately after the flare onset, the intercombination line increases dramatically while the forbidden line increases only somewhat. It is only later in the flare \( (t > 13 \text{ minutes } 39 \text{ s}) \) that the forbidden line gradually increases in intensity, an indication that the total amount of O^+6 ions also increases. The increase in the number of O^+5 ions is the result of ionization of O^+5 or recombination of O^+7 ions from the surrounding plasma regions. Figures 14a and 14b present the variations of density versus time in the \( 2 \times 10^6 \) K plasmas during the 1980 April 8 and May 9 flares as determined from the O vii line ratios. Most likely, these dramatic electron density increases result from plasma compression. Such a compression presumably can result from large electric currents.

3.3. The Morphology of \( 20 \times 20 \text{ K} \) Plasma in Large Nonimpulsive Solar Flares

The soft X-ray telescope (SXT) on the Japanese spacecraft Yohkoh (Tsuneta et al. 1991) was designed to record images of large solar flares. In a recent study of SXT results, we (Acton et al. 1992) have tried to confirm results obtained from the 1973 June 15 observations of the well-recorded large Skylab flare (Widing & Dere 1977; Cheng 1977). The Skylab observations revealed for the first time that the hot and dense flaring plasmas are located at loop tops while the loop legs are colder. Similar results were also reported from experiments on the Japanese satellite Hitori (i.e., Tanaka 1987). During the SXT studies we have identified 10 flares whose size is significantly larger than the instrumental spatial resolution (2.5 ≈ 2000 km). The SXT observations showed that within the instrument spatial resolution a substantial population of large nonimpulsive flares consist of single or multiple loop structures. The number of flaring loops is small (most often only one loop appears to exist), and seldom is the number of loops larger than three. In each of these flaring loops the bulk of the \( 2 \times 10^6 \) K radiation is emitted from compact regions composed of dense material \( (N_e \approx 10^{12} \text{ cm}^{-3}) \) at loop tops. The loop tops brighten very early in the event, perhaps simultaneously with flare onset, and stay bright well into the decay phase (Fig. 15 [Pls. 7 and 8]). The hot bright region often appears to consist mainly of a single SXT pixel \((1800 \times 1800 \text{ km}^2)\); it stays compact and does not spread down the legs of the loop(s). In a few cases, what appears to be loop footpoints brighten very early in the event. However, later in the flare, emission from the footpoints is significantly fainter than emission from the loop itself. A comparison between the lower temperature radiation transmitted by the thin Al filter and the higher temperature radiation transmitted by the Be filter (Tsuneta et al. 1991) suggests that the lower parts of flaring loops are significantly colder than loop tops. No large motions of hot plasmas from the chromosphere toward the loop tops are seen to account for the large brightness at the loop tops (Feldman 1990; Doschek et al. 1992). The high-density plasma responsible for the \( 2 \times 10^6 \) K emission, quite clearly, is heated and compressed in situ. It should be kept in mind that the SXT observations clearly confirmed the Skylab 1973 June 15 and Hitori results mentioned above, i.e., that the compact and round image in Fe xxiv \( (T_e \approx 2 \times 10^7 \text{ K}) \) arises from a loop(s) top, while the elongated Ca xvii flare image \( (T_e < 1 \times 10^7 \text{ K}) \) is shaped much like a whole loop in projection.

How could a blob of plasma with a temperature of \( T_e \approx 2 \times 10^7 \text{ K} \) and a density of \( N_e \geq 1 \times 10^{22} \text{ cm}^{-3} \), formed at the top of a loop, survive in that configuration for many minutes and in some instances for over an hour without spreading...
Fig. 15.—Images of a large limb flare recorded through a 119 μm Be filter by the soft X-ray telescope on Yohkoh (Tsuneta et al. 1991). (a) Flare images recorded every 12–24 s. (b) Enlarged images during the beginning, near maximum, and near the end of the same flare. The images are normalized to the brightest pixel in each image.

Feldman (see 411, 904)
FIG. 15b

FELDMAN (see 411, 904)
along the magnetic field lines? It is most reasonable to assume that the hot and dense blob is continuously being regenerated by a repeated string of large electric current instabilities similar to the ones observed in laboratory Z pinch devices (Burkhalter et al. 1979).

4. CALCULATIONS OF ELECTRICAL CONDUCTIVITY

The electrical conductivity of a mildly ionized gas can be evaluated under a variety of plasma conditions. When $E \parallel H$ the electrical conductivity ($\sigma$) of a mildly ionized gas is given by

$$\sigma = N_e e^2 \left( \frac{1}{m_e v_{en}} + \frac{1}{m_i v_{in}} \right),$$

where $v_{en}$ is the electron-neutral collision frequency and $v_{in}$ is the ion-neutral collision frequency. (The units used are in cgs, and the conductivity is in s$^{-1}$.) In the lower solar atmosphere electron collisions with H are the dominant process. At $T_e = 3000$ K we obtain that $v_{en} = 1.5 \times 10^{-7} N_H$. (The ion-collision term is small and can be neglected.)

By evaluating equation (1) we get $\sigma = 2 \times 10^{13} \times N_e/N_H + N_e$ s$^{-1}$. Therefore if Mg, Si, and Fe are ionized ($N_e/N_H + N_e = 1 \times 10^{-11}$) $\sigma = 2 \times 10^{11}$ s$^{-1}$. However if only K is ionized ($N_e/N_H + N_e = 1.3 \times 10^{-7}$), $\sigma = 3 \times 10^8$ s$^{-1}$. In cases where K is partially ionized or many of the free electrons become attached to neutrals and form negative ions, the electrical conductivity of the region will be significantly reduced.
The intensity ratio of the forbidden to the intercombination line for O vii as a function of the electron density at a $2 \times 10^6$ K temperature (Doschek et al. 1981).

The electrical conductivity of a mildly ionized gas where $E \perp H$ is given by

$$\sigma = e^2 \sum \frac{N_i v_{i\beta}/m_i}{\lambda_i + \omega_i} + \frac{N_e v_{e\alpha}/m_e}{\omega_e},$$

where $\omega_i = eB/m_i c$ and $\omega_e = eB/m_e c$. By substituting in the equation $N_H = 1 \times 10^{12}$ and $N_e/N_H+N_e = 1.3 \times 10^{-7}$, the electrical conductivity becomes $\sigma = 2 \times 10^{-1}$ s$^{-1}$. (For details see Akasofu & Chapman 1972, p. 131 and 240.) It is important to remember that the perpendicular conductivity is small not because of resistive dissipation, i.e., because magnetic energy is being converted into heat, but because the conduction electrons are constrained to circle the magnetic field.

Notice that the conductivity of a fully ionized gas at a temperature $T_e = 3000$ K is $\sigma = 3 \times 10^{12}$ s$^{-1}$ while the metallic conductivity is $\sigma \approx 1 \times 10^{16}$ s$^{-1}$ (Jackson 1973, p. 459).

As seen from the above, the conductivity of a mildly ionized gas changes dramatically with the ionization fraction of the gas and the strength of the magnetic field. When including the chemistry of the solar gas under decreasing temperature conditions the relative number of ionized particles may become extremely small, either because of the low ionization level or because of the formation of increasing numbers of molecules.

Oster (1968), Kopecký & Obržík (1968), and others attempted to calculate the conductivity in different regions of the solar atmosphere. They have estimated that in sunspot regions the electrical conductivity can be as low as $1 \times 10^9$ s$^{-1}$, suggesting that under such conditions the time scale for joule dissipation is similar to or perhaps smaller than the lifetime of the particular features.

5. MEASURING THE CONDUCTIVITY OF A SOLAR-TYPE ATMOSPHERE IN THE SUN AND IN THE LABORATORY

Parker (1979, p. 217) discusses a method by which the actual solar electric conductivity can be evaluated provided very high spatial resolution is available. The idea is based on the fact that the reduced electrical conductivity ($\sigma$) of the solar photosphere will lead to an increase of the diffusion velocity across the magnetic fields. The diffusion velocity is defined as

$$v = \frac{c^2}{\pi d \sigma},$$

where $c$ is the speed of light and $d$ is a characteristic size of the...
magnetic structures. In the case of fibrils where it is estimated that \( d \approx 200 \) km and for a conductivity \( \sigma = 3 \times 10^9 \) s\(^{-1}\), the velocity across the magnetic field should be \( v = 0.5 \) km s\(^{-1}\). Although a measurement of such a velocity under normal conditions is feasible, unfortunately we do not yet have the ability to resolve the individual fibrils and measure the velocity of the gas streaming across their field.

The knowledge of electrical conductivity properties of cold \((T_e \leq 3000 \) K\) solar and stellar type atmospheres is of fundamental importance for the physics of stellar atmospheres. It seems that it should not be an insurmountable task to measure in the laboratory the conductivity of an atmosphere containing all the important photospheric elements. After all, temperatures of \( T \leq 3000 \) K are being produced in the laboratory and the number of particles per unit volume required for such an experiment are on the order of \( 1 \times 10^{18} \) cm\(^{-3}\), within an order of magnitude from the number of particles in the Earth atmosphere. Elements having photospheric abundance ratios of \( N_e/(N_{\text{He}} + N_{\text{He}}) > 1 \times 10^{-5} \) can be included. It is conceivable that if the electrical properties of such a mixture are measured under a variety of temperature, density, and radiation field conditions appropriate for the coldest stellar regions in G, K, and M class atmospheres, more realistic atmosphere models will emerge.

6. CONCLUSION

The discussion in § 2 demonstrates that, because of the manner by which the abundance fraction versus FIP of elements in the solar photosphere are arranged, the electrical conductivity in the solar lower atmosphere is an extremely sensitive function of temperature. In cases where the temperature is sufficiently low, only atoms with FIP < 5 eV may be fully or perhaps only slightly ionized, implying an electron to neutral H + He atom abundance \( N_e/N_{\text{He}} \leq 1 \times 10^{-7} \). The number of free electrons can be further reduced by becoming attached to neutrals and forming negative ions or by forming molecules. The electrical conductivity of such plasmas may be sufficiently low to allow the buildup of large electrostatic fields. Joule energy dissipation in such regions can be responsible for the phenomena seen in the upper solar atmosphere. It was also suggested that the entire process of elemental abundance enrichment of the solar upper atmosphere can be associated with the presence of electrostatic fields.

Section 3 alludes to the fact that dramatic density increases are common occurrences in the solar upper atmosphere, a process that can be explained by large flows of electric current. The purpose of this paper was not to evaluate the electrical conductivity of cold regions in the solar atmosphere. However, for the sake of completeness § 4 presents a few calculations which show that changes in the conditions of the cold solar gas may cause large changes in the electrical conductivity.

It is an established belief among solar physicists that the solar atmosphere has sufficiently high electrical conductivity to prevent the generation of significant electric fields in all parts of the solar atmosphere. In this paper I have presented a number of observations which compel us to envision the existence of...
lower solar atmosphere regions sustaining strong electric fields. If indeed this is the case, then the electric fields may cause current flows which may account for the high temperature and the large electron densities detected in the upper solar atmosphere. Because of the difficulties and the fragmentary nature of solar observations, one may conceive different explanations to some of the phenomena presented. However, it is important to note that all the facts taken together present a compelling case for questioning the assumption that "Everywhere in the photosphere-corona domain the electric conductivity is HIGH." If indeed the conductivity in parts of the solar atmosphere is low, the high temperature in the corona and in Edlén's laboratory could have been produced by similar processes.

It is also interesting to note that the activity of solar-like stars seem to increase with the decrease in their surface temperature. Flare stars are M class stars, while hotter stars (F stars) with surface temperature higher than the Sun show a diminishing level of activity. Perhaps, as their surface temperature is reduced, their conductivity diminishes and larger parts of the star's surface can act as increasingly larger capacitors where energy for a variety of flare and coronal type phenomena can be stored. A similar effect is seen in the Sun. It is a well-known fact that the most active regions on the Sun, the sunspots, are among the coldest regions on the solar surface.

It is suggested that an attempt be made to measure in the laboratory, as a function of temperature, density, and radiation field, the conductivity of a gas with photospheric composition.

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