Stellar Atmospheres

Peter H. Hauschildt

Dept. of Physics and Astronomy & Center for Simulational Physics, University of Georgia, Athens, GA 30602-2451

France Allard

C.R.A.L (UML 5574) Ecole Normale Superieure, 69364 Lyon Cedex 7, France

Jason Aufdenberg

Dept. of Physics and Astronomy, Arizona State University, Tempe, AZ 852870-2504

Travis Barman & Andreas Schweitzer

Dept. of Physics and Astronomy, University of Georgia, Athens, GA 30602-2451

E. Baron

Dept. of Physics and Astronomy, University of Oklahoma 440 W. Brooks, Rm 131, Norman, OK 73019-0225

Abstract. We give an overview about the state-of-the-art in stellar (and substellar) atmosphere simulations. Recent developments in numerical methods and parallel supercomputers, as well as advances in the quality of input data such as atomic and molecular line lists have led to substantial improvements in the quality of synthetic spectra when compared to multi-wavelength observations. A wide range of objects from giants to main-sequence stars down to substellar objects is considered. We discuss effects such as limb darkening, atomic and molecular NLTE (and) line blanketing, winds, external irradiation, formation and opacities of dust particles and clouds; each of these affects the structure of the atmospheres and their spectra. Current models can simultaneously fit many of the observed features of a given star with a single model atmosphere. However, a number of problems remain unsolved and will have to be addressed in the future, in particular for very low mass stars and substellar objects.

1. Introduction

Stellar atmosphere modeling has experienced a renaissance in the past decade with the advent of better algorithms and faster computers. This has allowed...
research groups to remove or relax many of the “standard” assumptions that were made in the 70s through 80s and that had become accepted wisdom over the years. Surprisingly (or not) the new calculations show that many of these assumptions are actually quite bad and can lead to spurious results or incorrect interpretations of observed spectra. The intricate connection between geometry (plane parallel or spherical), line blanketing (atomic and/or molecular) and non-LTE effects (using small to extremely large model atoms and molecules) began to emerge slowly as crucial ingredients for physically correct and meaningful interpretations and analyses of stellar spectra. Unfortunately, easy and simple solutions do not really work for stellar atmospheres (although everybody likes the easy way out and some of them are useful for teaching purposes) and have actually hindered progress and reduced the reliability of results.

Table 1. Selected molecules considered in the EOS

<table>
<thead>
<tr>
<th>NH</th>
<th>C₂</th>
<th>CN</th>
<th>CO</th>
<th>MgH</th>
<th>CaH</th>
<th>SiH</th>
<th>TiO</th>
<th>H₂O</th>
<th>H₂</th>
</tr>
</thead>
<tbody>
<tr>
<td>N₂</td>
<td>NO</td>
<td>CO₂</td>
<td>O₂</td>
<td>ZrO</td>
<td>VO</td>
<td>MgS</td>
<td>SiO</td>
<td>AlH</td>
<td>HCl</td>
</tr>
<tr>
<td>HF</td>
<td>HS</td>
<td>TiF</td>
<td>AI</td>
<td>BO</td>
<td>C₄</td>
<td>NaO</td>
<td>MgO</td>
<td>ScO</td>
<td>YO</td>
</tr>
<tr>
<td>SIF</td>
<td>NaCl</td>
<td>CaOH</td>
<td>HCN</td>
<td>C₂H₂</td>
<td>CH₄</td>
<td>CH₂</td>
<td>C₂H</td>
<td>HCO</td>
<td>NH₂</td>
</tr>
<tr>
<td>LiOH</td>
<td>CaO</td>
<td>AIOD</td>
<td>NaOH</td>
<td>Mg₂H</td>
<td>AIOD</td>
<td>AIOD</td>
<td>AIOD</td>
<td>SiH₂</td>
<td>SO₂</td>
</tr>
<tr>
<td>H₂S</td>
<td>OCS</td>
<td>KOH</td>
<td>TiO₂</td>
<td>TiOCl</td>
<td>VO₂</td>
<td>FeF₂</td>
<td>YO₂</td>
<td>ZrO₂</td>
<td>BaOH</td>
</tr>
<tr>
<td>La₂O</td>
<td>C₄H₆</td>
<td>C₃</td>
<td>Si₂</td>
<td>AI</td>
<td>CH₃</td>
<td>Al₂</td>
<td>FeO</td>
<td>Sc</td>
<td>TiF₂</td>
</tr>
<tr>
<td>Al₂O₂</td>
<td>SiCl</td>
<td>C₃Cl</td>
<td>AI</td>
<td>FeF</td>
<td>C₂H</td>
<td>Al₂</td>
<td>FeO</td>
<td>Sc</td>
<td>TiF₂</td>
</tr>
<tr>
<td>KCl</td>
<td>CaCl</td>
<td>TiS</td>
<td>TiCl</td>
<td>SiN</td>
<td>AIF</td>
<td>AIF</td>
<td>AIF</td>
<td>C₂H</td>
<td>C₂H</td>
</tr>
<tr>
<td>Fe</td>
<td>LiCl</td>
<td>N₂</td>
<td>NH</td>
<td>SO</td>
<td>S</td>
<td>Al₂</td>
<td>AIF</td>
<td>C₂H</td>
<td>C₂H</td>
</tr>
<tr>
<td>AIO₂F</td>
<td>Al₂O₂</td>
<td>Al₂O₂</td>
<td>BO</td>
<td>BO</td>
<td>BO</td>
<td>BO</td>
<td>BO</td>
<td>BO</td>
<td>BO</td>
</tr>
<tr>
<td>BH₃</td>
<td>H₂O₂</td>
<td>BO</td>
<td>BO</td>
<td>BO</td>
<td>BO</td>
<td>BO</td>
<td>BO</td>
<td>BO</td>
<td>BO</td>
</tr>
<tr>
<td>BeC₂</td>
<td>Be₂</td>
<td>Be₂</td>
<td>BeH₂</td>
<td>BeH₂</td>
<td>BeO</td>
<td>BeO</td>
<td>BeO</td>
<td>BeO</td>
<td>BeO</td>
</tr>
<tr>
<td>CH₃</td>
<td>CH₂</td>
<td>CH₂</td>
<td>CH₂</td>
<td>CH₂</td>
<td>CH₂</td>
<td>CH₂</td>
<td>CH₂</td>
<td>CH₂</td>
<td>CH₂</td>
</tr>
<tr>
<td>CaO₂H₂</td>
<td>MgCl</td>
<td>SiCl</td>
<td>FeCl₂</td>
<td>C₂H₂</td>
<td>MgCl₂</td>
<td>MgCl₂</td>
<td>MgCl₂</td>
<td>TiOCl</td>
<td>ScCl₂</td>
</tr>
<tr>
<td>ZrCl₂</td>
<td>TiCl₄</td>
<td>ZrCl₄</td>
<td>ZrCl₄</td>
<td>C₂H₂</td>
<td>SiH₄</td>
<td>SiH₄</td>
<td>SiH₄</td>
<td>SiH₄</td>
<td>SiH₄</td>
</tr>
<tr>
<td>ZrF₂</td>
<td>TiF₃</td>
<td>ZrF₄</td>
<td>ZrF₄</td>
<td>FeO₂H₂</td>
<td>SO₄</td>
<td>(KOH)₂</td>
<td>(LH)O₂</td>
<td>Mg₂H₂</td>
<td>(NaOH)₂</td>
</tr>
<tr>
<td>PhH₂</td>
<td>PhH₂</td>
<td>SiH₄</td>
<td>SiN</td>
<td>PO₂</td>
<td>SO₂</td>
<td>PO₂</td>
<td>PO₂</td>
<td>PO₂</td>
<td>PO₂</td>
</tr>
<tr>
<td>H₃C₂N</td>
<td>C₂H₄N</td>
<td>C₂H₄</td>
<td>C₂H₄</td>
<td>C₂H₄</td>
<td>C₂H₄</td>
<td>C₂H₄</td>
<td>C₂H₄</td>
<td>C₂H₄</td>
<td>C₂H₄</td>
</tr>
<tr>
<td>HC₅N</td>
<td>C₂H₄</td>
<td>C₂H₄</td>
<td>C₂H₄</td>
<td>C₂H₄</td>
<td>C₂H₄</td>
<td>C₂H₄</td>
<td>C₂H₄</td>
<td>C₂H₄</td>
<td>C₂H₄</td>
</tr>
<tr>
<td>HC₅N</td>
<td>OH</td>
<td>CH</td>
<td>CH</td>
<td>CH</td>
<td>CH</td>
<td>CH</td>
<td>CH</td>
<td>CH</td>
<td>CH</td>
</tr>
<tr>
<td>CS⁻</td>
<td>FeO⁻</td>
<td>BO⁻</td>
<td>AlO₂⁻</td>
<td>AI⁻</td>
<td>AlO₂⁻</td>
<td>AlO₂⁻</td>
<td>AI⁻</td>
<td>AlO₂⁻</td>
<td>AI⁻</td>
</tr>
<tr>
<td>TiO²⁺</td>
<td>ZrO²⁺</td>
<td>AlO²⁺</td>
<td>BaO⁺</td>
<td>H₂O⁺</td>
<td>CaO⁺</td>
<td>SO⁺</td>
<td>H₂O⁺</td>
<td>H₂O⁺</td>
<td>H₂O⁺</td>
</tr>
</tbody>
</table>

New observational techniques opened and continue to open up new areas of stellar atmosphere research. Most importantly these are the observations of very low mass stars and, after decades of searching, brown dwarfs and extrasolar giant planets. Modeling these objects requires sophisticated stellar atmosphere type modeling with complex equations of state and hundreds of millions of molecular spectral lines in order to even approximately reproduce the observed spectra. These new observations have prompted further evolution of stellar atmosphere modeling and helped rejuvenate the field in general.

In the following we will briefly introduce the numerical methods that modern stellar atmosphere research employs (there are also plenty of legacy applications and codes that are still widely used) and then discuss some results that are of interest in the context of this meeting.

2. Methods and Models

For our model calculations, we use our multi-purpose stellar atmosphere code PHOENIX (version 10.9 Hauschildt, Baron, & Allard, 1997; Baron & Hauschildt, 1998; Hauschildt, Allard, & Baron, 1999; Hauschildt et al., 1999; Hauschildt & Baron, 1999). Details of the numerical methods are given in the above references, so we do not repeat the description here.
Table 2. Selected liquid/dust species considered in the EOS

<table>
<thead>
<tr>
<th>Al/II</th>
<th>B/II</th>
<th>Ba/II</th>
<th>Be/II</th>
<th>Ca/I</th>
<th>Cr/I</th>
<th>Cu/I</th>
<th>Fe/I</th>
<th>K/I</th>
</tr>
</thead>
<tbody>
<tr>
<td>Li/I</td>
<td>Mg/I</td>
<td>Mn/I</td>
<td>Na/I</td>
<td>Nb/I</td>
<td>Ni/I</td>
<td>P/I</td>
<td>S/I</td>
<td>Si/I</td>
</tr>
<tr>
<td>Sr/I</td>
<td>VO/I</td>
<td>Ti/I</td>
<td>V/I</td>
<td>Zn/I</td>
<td>Zr/I</td>
<td>BeO/I</td>
<td>Cr/I</td>
<td>NbO/I</td>
</tr>
<tr>
<td>Sb/I</td>
<td>CO/O</td>
<td>Cl/O</td>
<td>Fe/I</td>
<td>MgO</td>
<td>NiO</td>
<td>SiO</td>
<td>NaO</td>
<td>P/O</td>
</tr>
<tr>
<td>O2V/II</td>
<td>Sb/II</td>
<td>Nb/II</td>
<td>Si/II</td>
<td>Cr2O3</td>
<td>SiO2</td>
<td>CO2</td>
<td>CO2</td>
<td>O3</td>
</tr>
<tr>
<td>Sb2H/II</td>
<td>H2O/S</td>
<td>Ba2K2O6</td>
<td>K2O</td>
<td>Al</td>
<td>B</td>
<td>Ba</td>
<td>Be</td>
<td>C</td>
</tr>
<tr>
<td>Ca</td>
<td>Co</td>
<td>Ca</td>
<td>Cu</td>
<td>Fe</td>
<td>Mg</td>
<td>Mn</td>
<td>Na</td>
<td></td>
</tr>
<tr>
<td>Ni</td>
<td>Ni</td>
<td>P</td>
<td>S</td>
<td>Si</td>
<td>S</td>
<td>Ti</td>
<td>V</td>
<td>Zn</td>
</tr>
<tr>
<td>Zr</td>
<td>MgO</td>
<td>FeS</td>
<td>CaO</td>
<td>CaS</td>
<td>MgS</td>
<td>TiN</td>
<td>AlN</td>
<td>NiS</td>
</tr>
<tr>
<td>MnS</td>
<td>TiO</td>
<td>VO</td>
<td>CuO</td>
<td>FeO</td>
<td>TiCl</td>
<td>SiC</td>
<td>Zn</td>
<td>H2O</td>
</tr>
<tr>
<td>Ti2O</td>
<td>ZrO2</td>
<td>SiO2</td>
<td>Fe2O</td>
<td>NiO2</td>
<td>MgO</td>
<td>N2</td>
<td>50</td>
<td>0/0</td>
</tr>
<tr>
<td>MgTiO3</td>
<td>Mg2O3</td>
<td>SiO3</td>
<td>MnO3</td>
<td>Na2SiO3</td>
<td>K2SiO3</td>
<td>Fe2O4</td>
<td>CaSiO3</td>
<td>MgAl2O4</td>
</tr>
<tr>
<td>ZrSiO4</td>
<td>CaSiO3</td>
<td>NaSiO3</td>
<td>Al2SiO3</td>
<td>Al2SiO3</td>
<td>Al2O3</td>
<td>MgO</td>
<td>CaMg2Si2O7</td>
<td>Ca2Al5S</td>
</tr>
<tr>
<td>Cr23C6</td>
<td>KA2SiO3</td>
<td>NaAlSiO</td>
<td>Al2SiO3</td>
<td>MgC2</td>
<td>MgC3</td>
<td>Al4C3</td>
<td>Cr7C3</td>
<td>Cr7C3</td>
</tr>
</tbody>
</table>

Table 3. Complete listing of PHOENIX (version 10) atomic NLTE species. The table entries are of the form N/L, where N is the number of NLTE levels and L the number of primary NLTE lines for each model atom.

<table>
<thead>
<tr>
<th>Element</th>
<th>N</th>
<th>L</th>
</tr>
</thead>
<tbody>
<tr>
<td>He</td>
<td>19/37</td>
<td></td>
</tr>
<tr>
<td>Li</td>
<td>57/333</td>
<td>55/124</td>
</tr>
<tr>
<td>C</td>
<td>228/1387</td>
<td>85/336</td>
</tr>
<tr>
<td>N</td>
<td>252/2313</td>
<td>152/1110</td>
</tr>
<tr>
<td>O</td>
<td>36/66</td>
<td>171/1304</td>
</tr>
<tr>
<td>Ne</td>
<td>26/57</td>
<td></td>
</tr>
<tr>
<td>Na</td>
<td>53/142</td>
<td>35/171</td>
</tr>
<tr>
<td>Mg</td>
<td>573/673</td>
<td>72/340</td>
</tr>
<tr>
<td>Al</td>
<td>111/250</td>
<td>188/1674</td>
</tr>
<tr>
<td>Si</td>
<td>329/1871</td>
<td>93/436</td>
</tr>
<tr>
<td>S</td>
<td>329/1871</td>
<td>88/760</td>
</tr>
<tr>
<td>K</td>
<td>146/439</td>
<td>84/444</td>
</tr>
<tr>
<td>Ca</td>
<td>194/1029</td>
<td>87/485</td>
</tr>
<tr>
<td>Ti</td>
<td>395/5279</td>
<td>204/2399</td>
</tr>
<tr>
<td>Fe</td>
<td>494/6903</td>
<td>617/1367</td>
</tr>
<tr>
<td>Co</td>
<td>316/4428</td>
<td>255/2725</td>
</tr>
<tr>
<td>Ni</td>
<td>153/1690</td>
<td>429/7445</td>
</tr>
<tr>
<td>Sr</td>
<td>52/74</td>
<td>32/90</td>
</tr>
<tr>
<td>Ca</td>
<td>38/75</td>
<td></td>
</tr>
<tr>
<td>Ba</td>
<td>76/114</td>
<td>51/121</td>
</tr>
</tbody>
</table>

One of the most important recent improvements of cool stellar atmosphere models is that new molecular line data have become available. This includes the addition of the HITRAN92 (Rothman et al., 1992) data base, first incorporated in our “Extended” model grid (Allard & Hauschildt, unpublished). For the molecules with lines available from other sources (e.g. water vapor and CO) it is usually better to use other sources in model calculations because the HITRAN92 lists are fairly incomplete (they have been prepared for the conditions of the Earth’s atmosphere) although the line data are of very high quality. The OH molecule is an exception because the vibrational bands present in the HITRAN92 database are not present in the Kurucz CD 15 (Kurucz, 1993) dataset listing electronic transitions of OH. Therefore, we use both HITRAN92 and CD 15 OH line data simultaneously in the model calculations.

The most recent CO line list was calculated by Goorvitch (Goorvitch & Chackerian, 1994a,b). This list is more accurate than older CO line data. How-
Figure 1. The spectral energy distribution of ε CMa from 350 Å to 25 μm compared with the 21750 K non-LTE model (N1) and the ATLAS9 21000 K model. The EUVE data are corrected for the interstellar neutral hydrogen column assuming a neutral hydrogen column density of $N(\text{H}_1) = 5 \times 10^{17}$ cm$^{-2}$.

However, we have kept the electronic CO transitions available on Kurucz’s CD 15. Similarly, we have replaced the Kurucz CN line list (CD 15) with a more recent calculation by Jørgensen (Jørgensen & Larsson, 1990). CN lines are comparatively weak in most of the models presented here. In addition, we include line lists for YO (Littleton, 1987) and ZrO (Littleton & Davis, 1985) as well as H$_2^+$ partition functions and lines (Neale & Tennyson, 1995; Neale, Miller, & Tennyson, 1996).

A long-standing problem with M dwarf models was that TiO line lists were incomplete at high temperatures. The use of “straight means” (AH95 models) helped to block flux which otherwise escapes between lines in the incomplete list. But these models also blocked too much flux in most cases, and were only appropriate in late-type M dwarfs when TiO bands are very strong. A more complete line list was needed to model stars from the onset of TiO formation to its gradual disappearance from the gas phase in brown dwarfs. The AMES-TiO list (Schwenke, 1998) largely resolved this issue. We found that the list provides more opacity in most bands and adequately suppresses flux between
Figure 2. A comparison of synthetic radio spectra from A-type supergiant, expanding model atmospheres which have different mass-loss rates but otherwise fixed parameters \( T_{\text{eff}} = 9500 \text{ K}, \log(g) = 1.5, v_{\infty} = 260 \text{ km s}^{-1}, \beta\text{-law velocity field with } \beta = 3.0 \). Mass-loss rates have units of \( M_\odot \text{ yr}^{-1} \).

bands (Allard, Hauschildt, & Schwenke, 2000). New, smaller oscillator strength values also play an important role by systematically assigning cooler models (at least for early type M dwarfs) to a given star, and in this way contribute to broader bands and lower inter-band flux as well. The combination of these effects helps to resolve most of the previously observed discrepancy between models and observations in the optical spectral energy distribution (SED) and photometry of M stars. Leinert et al. (1999) note, however, that flux excess remains substantial in the visual spectrum, suggesting some further incompleteness or \( f_{\text{el}} \) inaccuracies of the new TiO in the a-f system.

The absorption coefficient of water has been a problem for modelers of cool dwarf atmospheres for decades. In our Base model grid we used the straight means of the Ludwig (1971) water opacity tables. These tables are known to overestimate the water opacity significantly at higher (and more importantly for M dwarfs) temperatures (Schryber, Miller, & Tennyson, 1995), but they are accurate at lower temperatures. The currently available list of water lines of Miller & Tennyson (MT-H\(2\)O, Allard et al., 1994) is incomplete and therefore
some water opacity will be missing in models with $T_{\text{eff}} < 4000$ K. Comparison with high-resolution observations in spectral regions where the MT-H$_2$O list should be complete shows that its accuracy is high (Jones et al., 1995). For most cool star models presented here, we have replaced the UCL water vapor line-list (Miller et al., 1994; Schryber, Miller, & Tennyson, 1995) used in Hauschildt, Allard, & Baron (1999) with the AMES water line-list (Partridge & Schwenke, 1997, hereafter: AMES-H$_2$O). This list includes about 307 million lines of water vapor. The introduction of the AMES-H$_2$O opacities brings solid improvements of the near-infrared SED of late-type dwarfs, but fails, as the AH95 models did, to reproduce adequately the $J-K$ colors of hotter stars. Water vapor is a more important factor for the structure of the atmosphere than TiO because its overall opacity is larger and its lines are closer to the peak of the SED than the TiO bands, so the flux blocking effect of water vapor is more important for the temperature structure than that of TiO opacities for these low temperatures.
Figure 4. Best fit of Allard & Hauschildt (1995b) to the spectrum of the dM8e star VB 10 (Allard & Hauschildt, 1995a). The corresponding H- continuum obtained by neglecting molecular opacities only in the radiative transfer (long dot-dashed) reveals the magnitude of these opacities in a typical late-type M dwarf. The Planck distribution of the same T_{eff} is also shown for comparison. From Allard et al. (1997).

Schwenke and collaborators at NASA AMES are preparing a new dipole moment function for H_2O, which may change the high temperature high overtone water bands, and help resolve this discrepancy in the near future. Until a revised version of the AMES-H_2O line list becomes available, we will maintain two sets of model atmospheres for cool stars which allow the investigation of these issues: the AMES grid based on the new TiO and H_2O lists, and the AMES-MT grid which rely on the AMES-TiO and Schryber, Miller, & Tennyson (1995) line lists.

Our combined molecular line list includes about 550 million molecular lines. The lines are selected for every model from the master line list at the beginning of each model iteration to account for changes in the model structure (see below). Both atomic and molecular lines are treated with a direct opacity sampling method (dOS). We do not use pre-computed opacity sampling tables, but instead dynamically select the relevant LTE background lines from master line lists at the beginning of each iteration for every model and sum the contribution of every line within a search window to compute the total line opacity at arbitrary
Figure 5. Spectral distributions of emerging fluxes at the stellar surface for 3,000 K models with metallicities corresponding roughly to the solar neighborhood ([M/H] = 0.0), halo ([M/H] = −2.0), and Population III ([M/H] = −4.0) stars. A black-body of the same effective temperature (smooth curve) is shown for comparison. From Allard et al. (1997).

wavelength points. The latter feature is crucial in NLTE calculations in which the wavelength grid is both irregular and variable from iteration to iteration due to changes in the physical conditions. This approach also allows detailed and depth-dependent line profiles to be used during the iterations. Although the direct line treatment seems at first glance computationally prohibitive, it can lead to more accurate models. This is due to the fact that the line forming regions in cool stars and planets span a huge range in pressure and temperature so that the line wings form in very different layers than the line cores. Therefore, the physics of the line formation is best modeled by an approach that treats the variation of the line profile and the level excitation as accurately as possible. To make this method computationally more efficient, we employ modern numerical techniques, e.g., vectorized and parallelized block algorithms with high data locality (Hauschildt, Baron, & Allard, 1997), and we use high-end workstations or parallel supercomputers for the model calculations.
Figure 6. Overview over selected departure coefficients for a NLTE model with $T_{\text{eff}} = 4000$ K, $\log(g) = 0.0$, and solar abundances.

In the calculations presented in this contribution, we have included a constant statistical velocity field, $\xi = 2 \, \text{km} \, \text{s}^{-1}$, which is treated like a microturbulence. The choice of lines is dictated by whether they are stronger than a threshold $\Gamma \equiv \chi_l / \kappa_c = 10^{-4}$, where $\chi_l$ is the extinction coefficient of the line at the line center and $\kappa_c$ is the local b-f absorption coefficient (see Hauschildt, Allard, & Baron, 1999, for details of the line selection process). This typically leads to about $10 - 250 \times 10^6$ lines which are selected from master line lists. The profiles of these lines are assumed to be depth-dependent Voigt or Doppler profiles (for very weak lines). Details of the computation of the damping constants and the line profiles are given in Schweitzer et al. (1996). We have verified in test calculations that the details of the line profiles and the threshold $\Gamma$ do not have a significant effect on either the model structure or the synthetic spectra. In addition, we include about 2000 photo-ionization cross sections for atoms and ions (Mathisen, 1984; Verner & Yakovlev, 1995).

The equation of state (EOS) is an enlarged and enhanced version of the EOS used in AH95. We include about 500 species (atoms, ions and molecules) in the EOS, cf Table 1. This set of EOS species was determined in test calculations. The EOS calculations themselves follow the method discussed in AH95. For effective temperatures, $T_{\text{eff}} < 2500$ K, the formation of dust particles has
Figure 7. This plot shows the difference between synthetic spectra calculated for a model atmospheres assuming complete settling of all formed dust particles below the layers where spectrum forms ("cond" model) and for a model that assumes that the dust particles remain close to the layer in which they formed ("dusty" model) for $T_{\text{eff}} = 1700$ K, $\log(g) = 4.5$ and solar abundances.

to be considered in the EOS. In our models we allow for the formation (and dissolution) of a variety of grain species, see Table 2 for a selection.

The NLTE treatment of large model atoms or molecules such as $\text{H}_2\text{O}$ and $\text{TiO}$ which have several million transitions is a formidable problem which requires an efficient method for the numerical solution of the multi-level NLTE radiative transfer problem. Classical techniques, such as the complete linearization or the Equivalent Two Level Atom method, are computationally prohibitive for large model atoms and molecules. Currently, the operator splitting or approximate A-operator iteration (ALI) method (e.g., Cannon, 1973; Rybicki, 1972, 1984; Scharmer, 1984) seems to be the most effective way of treating complex NLTE radiative transfer and rate equation problems. Variants of the ALI method have been developed to handle complex model atoms, e.g. Anderson's multi-group scheme (Anderson, 1987, 1989) or extensions of the opacity distribution function method (Hubeny & Lanz, 1995). However, these methods have problems if line overlaps are complex or if the line opacity changes rapidly with optical depth,
a situation which occurs in cool stellar atmospheres. The ALI rate operator formalism, on the other hand, has been used successfully to treat very large model atoms such as Fe directly and efficiently (cf. Hauschildt & Baron, 1995; Hauschildt et al., 1996; Baron et al., 1996). It allows us to currently treat the species listed in Table 3 in direct NLTE.

3. Results

In the following sections we will give a few representative results that highlight important new developments in stellar atmospheres.

3.1. Hot Stars

Historically, the assumption of plane parallel geometry was thought to be a good approximation for main sequence stars and many models for giants and supergiants did not use spherical geometry. However, a closer look on the effects of spherical geometry coupled to line blanketing shows that the situation is far
Figure 9.  Example of limb darkening for a subgiant model (calculated using spherical geometry and spherically symmetric radiative transfer) with $T_{\text{eff}} = 5000$ K, $\log(g) = 2.5$ and solar abundances. The computed (normalized to the center of the stellar disk) monochromatic intensities at $\approx 5000\text{Å}$ are shown (+ symbols) and compared to linear (dotted line) and square root (dashed line) limb darkening laws.

more complex. As an example, we consider the EUV spectrum of the B2 II star $\epsilon$ CMa. Spectroscopy of $\epsilon$ CMa (HD 52089, HR 2618) (Cassinelli et al. (1995)) below the Lyman edge by the Extreme Ultraviolet Explorer (EUV) satellite is possible due to the extremely low neutral hydrogen column density toward this star. At a HIPPARCOS distance of $132^{+11}_{-9}$ pc along an exceptionally rarefied interstellar tunnel extending out from the Local Bubble (Welsh, 1989), $\epsilon$ CMa is attenuated by a neutral hydrogen column density of less than $5 \times 10^{17}$ cm$^{-2}$ (Gry, York, & Vidal-Madjar, 1985). Consequently, $\epsilon$ CMa is an extremely important star for its contribution to the local interstellar hydrogen ionization, producing more hydrogen ionizing flux than all nearby stars combined (Vallerga & Welsh, 1996). The large EUV flux from $\epsilon$ CMa affects the ionization state of the Local Cloud, the region of neutral hydrogen concentrated within a few parsecs of the Sun, in which the solar system is embedded (Bruhweiler (1996)). However, the EUV spectrum of $\epsilon$ CMa could not be fit using standard plane parallel model
atmospheres; the observed spectrum displays an apparent EUV excess compared to such models (Fig. 1).

We found significant differences in the strength of the predicted EUV flux between our line blanketed spherical and line blanketed plane-parallel models. A more realistic model treatment of early B giants with a spherical geometry and NLTE metal line blanketing results in the prediction of significantly larger EUV fluxes compared with plane-parallel models. This result appears to explain a large part of the reported discrepancy between the observed EUV flux of early B giants such as ε CMa and the EUV flux predicted by plane-parallel LTE and NLTE line blanketed model atmospheres. It is important to note that the relative extension (e.g., the ratio of the outer to the inner radius of the model atmosphere) is small. This effect on the spectrum is the result of a change in the temperature structure of the spherical models compared to the plane parallel models and only occurs when line blanketing (and thus backwarming) is included (see Aufdenberg et al., 1998, 1999, for details).

The mass-loss rates of OB supergiants are often approximated using a fully ionized, uniform, spherically symmetric mass flow model where the terminal velocity and radio flux density are measured quantities (Scuderi et al., 1998; Drake & Linsky, 1989). This model predicts a frequency dependence for the flux density of $S_\nu \propto \nu^{0.67}$. The winds of A-supergiants are, however, only partially ionized. Thermal radio emission from partially ionized circumstellar environments and winds has been studied by Simon et al. (1983). They show that the power-law radio spectrum varies from the fully ionized, optically thin case ($S_\nu \propto \nu^{0.6}$) to the optically thick, Rayleigh-Jeans case ($S_\nu \propto \nu^2$). The steepness of the radio spectrum depends on the degree of ionization and radial extent of the ionized envelope. As the extended envelope becomes less ionized, the spectral slope steepens and the radio flux drops. Our extended, expanding model atmospheres assume neither a constant temperature distribution nor constant ionization.

Figure 2 shows model IR-radio spectra of A-type supergiants with a range of mass-loss rates from zero to $10^{-6} \, M_\odot$ yr$^{-1}$. The frequency dependence of the synthetic radio spectra differ significantly from $S_\nu \propto \nu^{0.67}$. The static model shows the expected optically thick Rayleigh-Jeans spectrum ($S_\nu \propto \nu^2$) while the wind model spectra beyond $\sim 0.1$ cm follow $S_\nu \propto \nu^{1.1}$. The flux density longward of 0.01 cm increases significantly with $\dot{M}$ because the total density, including the ion density, is increased in the wind.

We also find that departures from LTE are quite important. Predicted flux densities in LTE are about an order of magnitude lower than corresponding non-LTE predictions with otherwise identical model parameters. In the non-LTE models hydrogen is ionized out to a radius ~4 times larger than in the LTE models. The temperature structure at the formation depths of the radio continuum is quite insensitive to $T_{\text{eff}}$. As a result, the radio flux densities predicted by the LTE models, where the ionization structures are determined purely by the local electron temperature, are also quite insensitive to $T_{\text{eff}}$. In contrast, temperature structure in the hydrostatic zones immediately below the wind is sensitive to $T_{\text{eff}}$ and in non-LTE, it is the Balmer continuum radiation field from these hydrostatic layers which controls the degree of hydrogen ionization in the outer wind.
For the A-type supergiant Deneb, IR spectrophotometry, the radio measurement at 3.6 cm, and upper limits at 2 cm and 6 cm are compared with two synthetic radio spectra in Figure 3. The non-LTE model, which is our best fitting model to the UV-optical-IR SED, predicts radio fluxes consistent with the upper limits and the 3.6 cm detection. While the LTE model has a very similar radio spectrum and provides a good fit to Deneb's UV-optical continuum, comparisons with the observed high-dispersion UV spectrum reveal that the non-LTE model is the overall better representation of Deneb's SED.

3.2. Cool Stars

Line blanketing  The number of molecular lines that are important in M-dwarf (and later) atmospheres are quite large. About 215 million molecular lines are selected (see above) for a typical giant model with $T_{\text{eff}} \approx 3000$ K whereas about 130 million molecular lines have to be considered for a dwarf model with the same effective temperature. The large “density” (in wavelength space) of molecular lines causes large line blanketing effects as illustrated in Figure 4 for a very simple case. The nearly complete coverage of the optical spectrum by TiO lines and of the near IR spectrum by H$_2$O lines effectively locks the peak of the spectral energy distribution in place at around 1.1$\mu$m even for substantially different $T_{\text{eff}}$, in stark contrast to the behavior expected for blackbodies. Line blanketing also produces a strong metallicity effect on the spectra as illustrated in Figure 5. Lowering the metal abundances reduces both the TiO and H$_2$O opacities by roughly the same amount. However, the H$_2$O opacity in the near IR is replaced by increasingly (with lower metallicity) stronger collision-induced opacities (due to the larger pressures in the spectrum forming regions). Therefore, the spectrum gets bluer with lower metallicities, even for comparatively low effective temperatures.

NLTE effects  Due to their very low electron temperatures, the electron density is extremely low in M stars; the absolute electron densities are even lower than found in low density atmospheres, such as those of novae and SNe. Collisional rates due to collisions with electrons, which tend to restore LTE, are thus very small in cool stars. This in turn could significantly increase the importance of NLTE effects in M stars when compared to, e.g., solar type stars with much higher electron densities and temperatures. Collisions with molecular hydrogen and helium will at least in part compensate for the diminished electron collisions, but cross-sections for these processes are not very well known. Therefore, the assumption of LTE for atoms and molecules in cool stars is by no means certain and needs to be verified for each species individually. Therefore, we have performed test calculations to place an upper limit on the importance of atomic NLTE effects in cool stars by only considering electron collisions.

We have calculated a small number of NLTE models in order to investigate the importance of NLTE effects on the structure of the model atmospheres. The results for cooler models were discussed in Hauschildt et al. (1997) and are not repeated here. Figure 6 shows an overview of selected NLTE species for models with $T_{\text{eff}} = 4000$ K, log($g$) = 0.0 and solar abundances. The total number of NLTE levels in each model is 4532 with a total of 47993 primary NLTE lines (see Hauschildt & Baron, 1999; Hauschildt et al., 1999, and references therein.
for details). The following species (and number of levels) were treated in NLTE: H I (30), Mg I (273), Mg II (72), Ca I (194), Ca II (87), Fe I (494), Fe II (617), O I (36), O II (171), Ti I (395), Ti II (204), C I (228), C II (85), N I (252), N II (152), Si I (329), Si II (93), S I (146), S II (84), Al I (111), Al II (188), K I (73), K II (22), Na I (53), and Na II (35). For most of the species, the departure coefficients are always close to unity, in particular for species with resonance lines and photoionization edges in the UV part of the spectrum. The species shown in Figure 6 are species with the most pronounced departures from LTE. The departures are generally too small to significantly affect the structure of the atmospheres. Results for NLTE calculations for the CO molecule show that the high cross-sections of H2 and He collisions restore LTE very successfully in the case of dwarfs stars (Schweitzer, Hauschildt, & Allard, 2000).

Dust and cloud formation The effects of dust formation on the atmosphere are mainly (a) the removal of important opacity sources (e.g., TiO, VO) from the gas phase and a corresponding weakening of their spectral lines; and (b) the presence of additional opacities produced by the grains themselves. The latter depends on the behavior of the macroscopic dust particles:

a) They might remain in the layers where the dust originally formed and thus cause strong optical and IR opacities (“Dusty” models).

b) They could rain out and settle below the line and continuum forming regions, in this case no grain opacities would be detectable in the spectrum (“Cond” models).

c) They can form clouds in the atmosphere so that dust opacities would only be present in the cloud layers but not in all the layers where the dust originally formed (“Cloudy” models).

Observational evidence suggests that case (a) is realized for Teff > 1700 K (late M dwarfs to early L dwarfs) whereas for both giant planets and extreme T dwarfs case (b) appears to be more appropriate. In the intermediate regime, clouds appear to form and case (c) is the best approximation. Figure 7 illustrates the general trend for an example model atmosphere. At the present time, this sequence is still very tentative and the physical models of dust and cloud formation have to be refined to obtain a truly physical picture of brown dwarfs and extrasolar giant planets.

Irradiation One feature of the extrasolar giant planets that have been discovered so far is that they are close to a more massive and hotter companion. Under such conditions, the structure of the atmosphere of the cooler companion (secondary) will be strongly affected by the presence of the strong radiation field of the hotter (primary) component. This will give rise to strong NLTE effects in the secondary’s atmosphere. If the separation of the two components is relatively large (but still small enough that the primary’s radiation field has an effect on the secondary), the calculation of an irradiated atmosphere is straightforward. If the separation is small, then the irradiated part of the secondary’s surface can be approximated by a patchwork of plane parallel surface elements. In the case of cool stars with strong convection in the lower atmosphere, the situation is somewhat more complex and entropy-matching procedures need to be
performed (Brett & Smith, 1993). Ultimately, the radiation transfer problem in the case of an irradiated atmosphere has to be solved using a full 3D approach. This is presently not feasible with the same complex input physics that can be included in 1D models, but simplified 3D models are possible and over the next decade or so computers should become powerful enough to handle the full 3D problem.

An important question concerns the input radiation field used in the irradiation modeling. A simple approximation would be to use blackbody radiation fields with a given color temperature. However, stellar radiation fields are very different from Planckian fields, in particular in the optical and UV spectral region where the differences can be orders of magnitude because of the very complex wavelength dependence of the stellar opacity. Blackbody radiation fields will only be useful for the simplest tests. Therefore, we calculate the input radiation fields with PHOENIX and thus have full control over the wavelength and angular grid on which the input field will be available. As an example, we show in Figure 8 the effects of irradiation on the emitted spectrum of a $T_{\text{eff}} = 500$ K dusty giant planet with a DM6 primary. Although this is a simple example of an irradiated atmosphere, it clearly shows the combined effects of reflection and heating of the outer atmosphere of the planet by the external radiation field.

3.3. Microlensing and Stars

One of the purposes for this meeting is to bring together different areas that contribute to our understanding of microlensing and what we can learn from it. From the stellar atmosphere point of view there are (at least!) two important areas that can be addressed by microlensing observations: limb darkening and star spots. Limb darkening is an important tool for the investigation of the structure of stellar atmospheres, in particular for subgiant and giants. In Figure 9 we show the limb darkening in the optical ($\sim 5000\AA$) for a cool giant with $T_{\text{eff}} = 5000$ K and $\log(g) = 2.5$ (solar abundances). For comparison, linear and square-root limb darkening laws are also shown. The form of the limb darkening and the deviations of the computed intensities from the simple limb darkening laws changes substantially with wavelength and model parameters (see also Orosz & Hauschildt, 2000). This is caused by the effects of spherical radiation transport, which tends to concentrate the emitted intensities closer to the center of the stellar disk than the plane parallel approximation. This is due to the fact that curvature close to the rim of the stellar disk actually reduces the total emissivity and optical depth and thus produces less emitted radiation than plane parallel models in which the pathlength and optical depth close to the rim actually approach infinity. This illustrates that in order to obtain reliable and useful stellar atmosphere information from microlensing observations we need to use rather sophisticated modeling; simple approximation will in most cases not work. Using detailed limb darkening models as shown here has helped Orosz & Hauschildt (2000) to address a number of problems regarding lightcurves of eclipsing binaries. The effects of star spots on microlensing observations is discussed in detail in the paper by P. Sackett (this volume).
4. Summary and Conclusions

In this paper we have discussed a few new results of stellar atmosphere modeling that have helped to resolve some outstanding problems with understanding and interpreting observed stellar spectra. During the last decade, progress was made by breakthroughs in both methodology and computer technology, which has led to substantially improved models and synthetic spectra. In many cases, even our current "best effort" models cannot reproduce observed spectra satisfactorily; this is in particular the case for L and T dwarfs. However, this is due to physical effects that we "know" but we cannot currently describe well enough (e.g. incomplete line lists for key molecules or dust and cloud formation processes). Another area that requires much more work is our detailed understanding of winds from both hot and cool stars. There is currently a lot of effort being put into the solution of these key problems, although it is clear that once they are solved, others will pop up in unexpected places.

Acknowledgments. This work was supported in part by the CNRS, INSU and by NSF grant AST-9720704, NASA ATP grant NAG 5-8425 and LTSA grant NAG 5-3619, as well as NASA/JPL grant 961582 to the University of Georgia, NASA LTSA grant NAG5-3435 and NASA EPSCoR grant NCCS-168 to Wichita State University. This work was supported in part by the Pôle Scientifique de Modélisation Numérique at ENS-Lyon. Some of the calculations presented in this paper were performed on the CNUSC IBM SP2, the IBM SP2 and SGI Origin 2000 of the UGA UCNS, on the IBM SP "Blue Horizon" of the San Diego Supercomputer Center (SDSC), with support from the National Science Foundation, and on the IBM SP and the Cray T3E of the NERSC with support from the DoE. We thank all these institutions for a generous allocation of computer time.

REFERENCES

Hauschildt et al.

Kurucz, R. L. 1993, Molecular data for opacity calculations, Kurucz CD-ROM No. 15.
Littleton, J. E. 1987, private communication.

Discussion

**Gray:** The red giant limb darkening law you displayed showed a very sharp drop towards the limb. How does this vary with different parameters? Is it sensitive? Does it keep the same sharp edge?

**Hauschildt:** The edge is very dependent on log g until reaching the dwarf regime (when the standard limb-darkening law is recovered). The dependence on $T_{\text{eff}}$ is less pronounced and much weaker. The location of the edge is wavelength dependent. The shape is slowly changing towards the dwarf regime.

**Zinnecker:** Can you summarise again how much the dust opacity affects the spectral appearance of M, L, T dwarfs and ultimately Jupiter-like planets?

**Hauschildt:** A very approximate and tentative picture is like this: Early M dwarfs: too hot for dust. Late M dwarfs ($T_{\text{eff}} < 2800K$): dust begins to form, small effect on spectra. “early” L dwarfs: dust opacity becomes strong; shallow H$_2$O bands by back-warming.

“Mid-late” L dwarfs ($T_{\text{eff}} < 2200K$): strong effect of dust opacities; “full dust”. Late L dwarfs ($T_{\text{eff}} = 1500K$?): Less deep opacity, possibly due to retraction of convection zone deeper into atmosphere; cloud formation.

T dwarfs, “Jovian” planets: even less dust opacity, but still important; formation of clouds; the atmospheres appear “clearing” with low $T_{\text{eff}}$.

**Gould:** Is spectral resolution of the atmosphere of any interest to you? Would it help resolve theoretical questions?

**Hauschildt:** Yes. Different lines, better spectral features, originate in different segments of the atmospheres. Extracting this information from microlensing events could help to constrain the structure of the atmosphere predicted by models.

**Sackett:** If you could choose, what wavelength range should be probed at what resolution for giants?
Hauschildt: There is no easy answer for this that I can think of. This can be determined by looking at the models (and the possible observations). I think that the NIR range around band heads and close to resonance lines is most likely to be useful. The actual “best range” will also be different for dwarfs and giants.

Kerins: In your spectral energy distribution fit to Kelu-1 most of the discrepancies with the real spectrum involve overestimates of the flux. However, just longward of 1μm the fit significantly underestimates the flux. Is there any significance to this?

Hauschildt: Yes, this is (mostly) due to known deficiencies of the molecular line data for H₂O (IR) and TiO. The latter have been addressed in new TiO line lists by Schwenke(1998) which effectively “kills” the excess peaks in the visible. New H₂O line lists are in preparation.

Gaudi: How well do your models agree with models of other groups?

Hauschildt: In general comparisons are hard due to the fact that groups use different input physics. For similar setups, results agree within reasonable limits, i.e. quite well.