

# NUCLEAR CHEMISTRY AND COSMOLOGY

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## I. INTRODUCTION

### What is Cosmology?

Cosmology is the study of the development of the universe, from its very "beginnings" to the present and into the future.

We try to gain an understanding of what has happened far from us, in space and in time, in terms of what we have learned in our development of science on Earth. Thus, we assume that the ideas of (mainly) physics and chemistry that we know to be valid "in the vicinity of the Earth, over the past few hundred years" — as expressed in the equations that we have developed to describe nature and the values we have obtained for the "fundamental constants"— are also valid at all times and all places throughout the universe.

Cosmology draws on results from a wide variety of scientific fields, such as astrophysics, astronomy, nuclear physics and chemistry, elementary particle physics, etc. This information has been used to develop scenarios, called "models," that have been successful in describing and explaining what is happening and what has happened in the universe, over a time scale of 15 billion ( $15 \times 10^9$ ) years!

### Nuclear Processes: the Key to the Universe.

Our knowledge of the characteristics of stars, such as their size, mass, age and luminosity (energy release), makes it clear that chemical reactions and other "ordinary" forms of energy cannot provide adequate power to the stars over their long lifetimes. Only the huge energy output from nuclear processes is sufficient for this purpose. This realization is one of the keystones of cosmology, and provides the reason that nuclear chemists are prominently involved in this field.

My main theme in this talk will be our observation of neutrinos as a probe of the nuclear reactions that occur in the sun. To put the discussion in perspective, I will spend some time discussing the beginnings of the universe and how we think stars are powered.

### Energy: Nuclear vs. Chemical.

What are the energies that characterize nuclear, in contrast to chemical, reactions? Nuclear reactions generally occur at million-electron-volt energies, mega-eV or MeV. A nuclear reaction that releases 1 MeV per nucleus releases 20 billion or  $2 \times 10^{10}$  calories per gram-atom (or mole, which contains  $6 \times 10^{23}$  nuclei), whereas a more familiar source of energy, the chemical reaction in which carbon burns in oxygen,  $C + O_2 \rightarrow CO_2$ , releases "only" 94 thousand (kilo) calories per mole or

4.1 eV per atom of carbon. Typical energy outputs of nuclear processes are  $10^5$  to  $10^6$  times larger than those in chemical reactions.

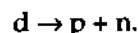
## II. SOME BASIC CONSIDERATIONS.

### Energies and Temperatures.

We will be discussing energy production in the universe. To do so, we will want to relate temperature ( $T$ , in degrees K) to energy ( $E$ ) via the equation,  $E = kT$ , where  $k$ , called Boltzmann's constant, has the value  $8.63 \times 10^{-5}$  electron volts (eV) per degree. Thus, 1 eV = 12,000 K (to two significant figures).

What is the significance of this result for cosmology? We must realize that different processes occur in different energy and temperature regimes. For example:

At energies of a few MeV, where 1 MeV corresponds to a temperature of 12 billion degrees ( $12 \times 10^9$  K), nuclei can readily combine in "fusion" reactions because their energies are greater than the electrical repulsion between the positively charged nuclei (called the "Coulomb barrier"). On the other hand, one of the first products of nuclear fusion, the deuteron ( $d$ ) is unstable at energies above 2.2 MeV. Above this "binding energy,"  $d$ , which is the isotope of hydrogen with mass 2 (also written as  ${}^2\text{H}$ ), breaks apart into a proton and a neutron,



At lower temperatures, around  $12 \times 10^6$  K, or 1 kilo-eV (keV), nuclear reactions can still occur in light elements, but relatively slowly. And, because the binding energies of the orbital electrons in light atoms are low, e.g., 14 and 25 eV respectively in hydrogen and helium, these elements are fully ionized at these temperatures. Note that the temperature at the center of the sun is just in this range, 14 million K, and decreases to about 6,000 K at the surface. The central temperature is still high enough for nuclear reactions to occur.

At much lower temperatures, the energies are just too low for nuclear reactions. "Chemistry" can begin at temperatures around 12,000 K, or 1 eV, with nuclei of light elements forming atoms by the addition of orbital electrons.

## III. THE BIG BANG SCENARIO.

### "In The Beginning..."

The so-called standard model considers that the universe had a "beginning" in which all of the available energy was confined to a small region of space, at

extremely high density and temperature. This energy was released in an "explosion," the Big Bang, which caused the universe to expand rapidly.

If we think of the early universe as being analogous to an expanding gas, we can understand what happened. Expansion of the gas caused it to cool, with a resultant decrease in temperature. However, at certain temperatures, "changes of state" occurred; just as a cooling gas reaches a characteristic temperature where the gas condenses to a liquid, so the temperature of the cooling universe reached a value where a particular elementary particle could be formed (remember the conversion of energy to mass via  $E = mc^2$ ).

In the model, after about 1 second had elapsed since "time zero" of the Big Bang, the following particles had already "frozen out": photons; "leptons," namely electrons and positrons (the antiparticle of the electron), and neutrinos and antineutrinos; and "nucleons" with atomic mass of one, namely protons ( ${}^1\text{H}$ , with atomic number  $Z = 1$  and atomic mass  $A = 1$ ) and neutrons. The temperature by this time had cooled to some  $10^{10}$  K, or MeV, which was still too hot for the effective formation of the deuteron,  ${}^2\text{H}$ , from  ${}^1\text{H} + n$ .

Then, as the universe continued to expand rapidly, a temperature was reached, in the  $10^9$  K range or hundreds of keV, where finally,  ${}^2\text{H}$  could be formed and survive. A sequence of exoergic nuclear fusion reactions could then occur, such as  ${}^2\text{H} + n \rightarrow {}^3\text{H}$  and  ${}^3\text{H} + {}^1\text{H} \rightarrow {}^4\text{He}$  ( $Z = 2$ ); or  ${}^2\text{H} + {}^1\text{H} \rightarrow {}^3\text{He}$  and  ${}^3\text{He} + n \rightarrow {}^4\text{He}$ ; followed by  ${}^3\text{He} + {}^4\text{He} \rightarrow {}^7\text{Be}$  ( $Z = 4$ ) and  ${}^7\text{Be} + e^- \rightarrow {}^7\text{Li}$  ( $Z = 3$ ) + neutrino (an electron-capture process). Stable nuclei heavier than  ${}^7\text{Li}$  (i.e., with  $Z > 3$  or  $A > 7$ ) are not formed, because the intermediate nuclei needed in the reaction sequence, with mass numbers of 5 and 8, are extremely unstable and don't last long enough to react any further.

The probabilities for these different reactions to occur can be measured by scientists in nuclear physics experiments. Incorporated into the Big Bang model, they are used to predict the relative importance of the different reaction paths, and the abundances of their end products in the universe.

We see that nuclei, with atomic numbers up to 4, were formed in these fusion reactions that began with  ${}^1\text{H} + n$ . However, in the model, as the universe continued to expand and cool, lower temperatures were attained where the nuclear processes could no longer occur. Element production ceased about three minutes after the Big Bang. Any free neutrons that remained then decayed with a half-life of 10.4 minutes, via the reaction  $n \rightarrow p + e^- + \bar{\nu}$ , the antineutrino (pronounced "nu-bar"). These residual neutrons were all converted to  ${}^1\text{H}$ .

So the universe, at the end of this period of "primordial nucleosynthesis," was composed of mainly the elements hydrogen and helium, in particular  ${}^1\text{H}$  and  ${}^4\text{He}$ . Formation of heavier elements would have to wait for other processes to occur.

Before we discuss some of these further developments, let's pause to see what evidence exists in support of the Big Bang model. After all, science is based on experimental data. If theory does not fit the facts, then theory has to be changed.

#### Data In Support of the Big Bang.

Several key observations [and their interpretations] led to the development of the Big Bang scenario.

- Wavelengths of emission spectra from galaxies are shifted "to the red." [These observed shifts, which are similar to the well-known Doppler shift that alters the frequency of emitted sound waves from a moving object, indicate that the galaxies are receding from us, and that the farther the galaxy, the faster it is moving. The universe is expanding from a common source: the rapid expansion tells us that the universe was once much hotter than it is today; also, the distances to the farthest galaxies allow us to estimate the age of the universe as  $15 \times 10^9$  years.]
- There is a uniform background of thermal radiation (energy) in space, equivalent to an ambient temperature of 2.8 K. [This temperature is taken to be the remnant of energy that remains today from the initial Big Bang.]
- Hydrogen and helium account for about 98% of all of the chemical elements in the universe, with a ratio of  ${}^4\text{He}$  to  ${}^1\text{H}$  of 7.7%. [These elements were the earliest ones that were formed. Their abundances can be predicted by the model.]
- The concentration of  ${}^4\text{He}$  is uniform throughout the universe. [This uniformity implies that the helium was produced in a short period of time, in a relatively small region of space. Otherwise, we would expect the concentration of helium to vary considerably in stars of different ages.]

#### IV. WHAT NEXT?

##### Nuclear Processes in Stars.

We have already learned two key points about our universe:

1) Nuclear reactions provide the mechanisms to produce the chemical elements, starting with the "building blocks" provided by the Big Bang. 2) These same nuclear reactions release energy, at least in the lightest elements. How economical! Nuclear processes "kill two birds with one stone."

However, the model requires that nuclear reactions stopped shortly after the Big Bang, because the expanding universe had become too cold. So, how were elements heavier than helium ever produced?

The conventional wisdom is that the "spark" that reignited these reactions was provided by the force of gravity, which led to the buildup of large local concentrations of matter at various points in space. The "seeds" for these concentrations, which were to develop into stars and galaxies, were probably just the result of random fluctuations in the density of the matter in the

expanding universe. Putting more matter into a given volume of space is equivalent to a compression. Just as an expanding gas cools, so a compressed gas heats up. Compression due to gravity thus provides a mechanism for reheating these regions of localized matter.

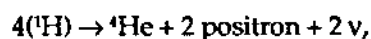
Eventually, temperatures around  $10^7$  K, or keV energies, are reached in the cores of stars, where nuclear reactions can again begin with hydrogen. However, the sequence of reactions, in which hydrogen fuses to form helium and release energy in the keV range, is very different from the sequence described earlier, which took place under much hotter conditions, around 1 MeV. As we will see, these differences are extremely important for the subsequent development of the universe.

At these keV energies,  $^2\text{H}$  is stable, but no free neutrons are available to react with protons to form  $^2\text{H}$ . Instead, a very different route must be followed: two protons must fuse in a reaction that converts a proton into a neutron,



This "pp reaction" is a form of beta decay, which is controlled by the "weak" nuclear force. The rate of this reaction is extremely slow. It has been estimated that, in the center of the sun, the characteristic time for the fusion of two protons to form a deuteron is about  $10^{10}$  years! The low energies at which these fusion reactions occur and their very low rates ensure that stars live for billions of years. At higher temperatures, stars consume their nuclear fuel much faster.

The sequence of reactions that begins with this fusion of two protons is called "hydrogen burning." It provides the major pathway for energy production in main sequence stars, such as our sun. The reaction chain continues with  $^2\text{H} + ^1\text{H} \rightarrow ^3\text{He}$ , and then  $2(^3\text{He}) \rightarrow ^4\text{He} + 2(^1\text{H})$ . Note the absence of any reactions involving free neutrons. The overall reaction is



in which 26.7 MeV are released. This energy is produced at the star's core and radiated away at the surface. A balance is struck in the star between the expansion due to nuclear energy production and the compression due to gravity. The star's structure is stabilized by these competing forces, so long as its supply of hydrogen fuel lasts.

We have finally arrived at the point where we can talk about the detection of neutrinos from the sun!

#### What about Elements Heavier than Helium?

But first, let's say a few words about the formation of the rest of the chemical elements. I won't have time to go into much detail, and instead I refer you to the article by Professor Vic Viola that will soon appear in

the Journal of Chemical Education. The main point is that all of the chemical elements are formed in nuclear processes, most of which occur under conditions of high temperature and high density in stars.

To produce elements heavier than helium in any significant amounts requires higher temperatures and densities than occur in main sequence stars. For example, helium can "burn" in large stars called red giants, producing elements such as  $^{12}\text{C}$  and  $^{16}\text{O}$ . These nuclei can then combine to form even heavier products.

An interesting point here is that a variety of nuclear reactions that release energy can continue to occur, until the isotope  $^{56}\text{Fe}$  is reached. Because (as we will see shortly) this nuclide is the most stable isotope in nature, it is not possible to add any particles to  $^{56}\text{Fe}$  without providing energy as well (endoergic reactions). A major change occurs at this isotope; new types of nuclear processes must be "turned on" to produce the elements above iron. Some of these processes have to occur at extremely high temperatures and densities, such as those that occur in catastrophic supernovae explosions. The Viola article discusses these items in detail.

#### Understanding Elemental Abundances.

Without delving any deeper into the ways that these nuclear reactions occur, we can get some general understanding of what happens by looking at the end products of these nuclear processes, the chemical elements that exist in our solar system. Data on the abundances of the elements are shown in Figures 1 and 2, which were kindly provided by Professor Viola. A few trends are readily apparent.

Figure 1 shows the abundances of the elements plotted vs. atomic number up to  $Z = 30$ . After the very

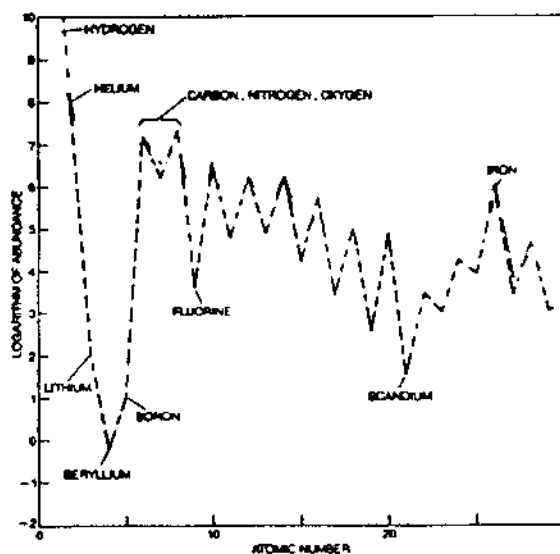


Figure 1. Relative abundances of the lighter chemical elements in the solar system, plotted vs. atomic number (figure was provided by V. Viola).

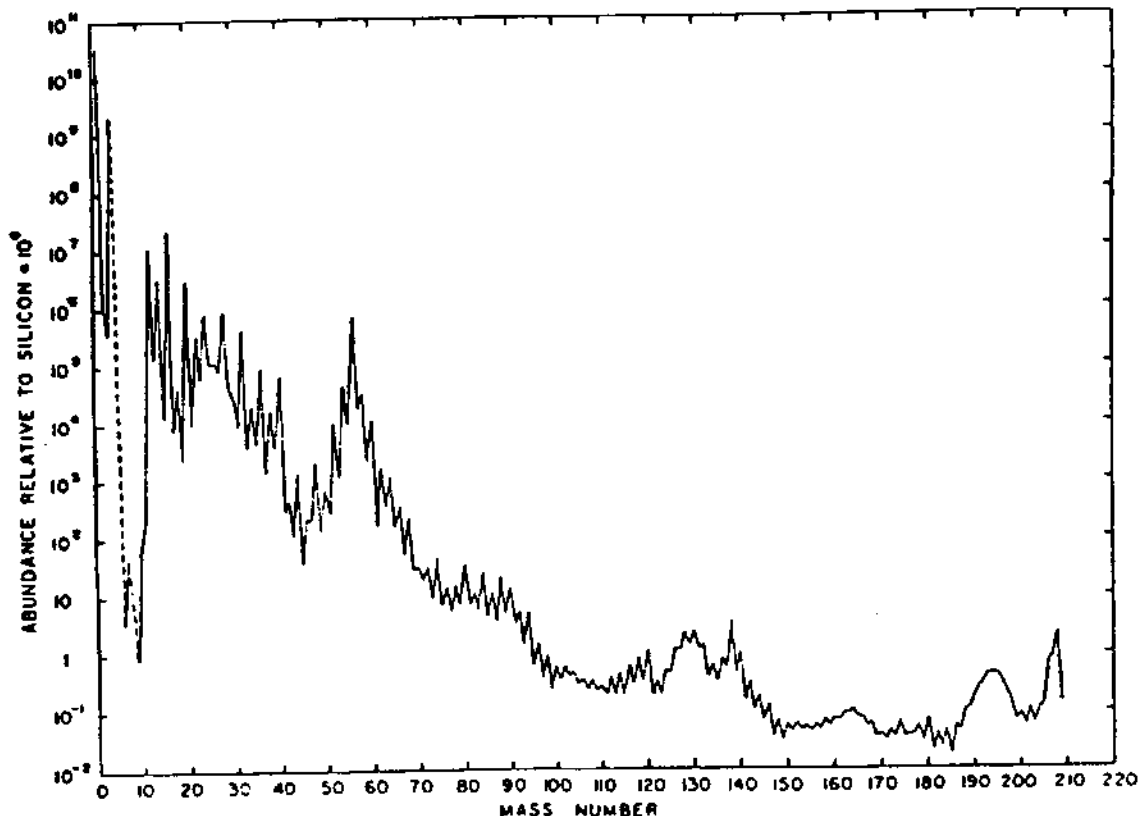


Figure 2. Abundances of the chemical elements in the solar system, plotted vs. atomic mass. Vertical scale is defined with silicon set at  $10^6$  atoms (figure was provided by V. Viola).

low results for Li, Be and B ( $Z$  from 3 to 5), the abundances from C to Ca ( $Z = 6$  to 20) show a regular decrease, on which is superimposed a pronounced saw-tooth pattern, with the even- $Z$  elements being higher than their odd- $Z$  neighbors. The fact that the abundances then rise to a peak at  $^{56}\text{Fe}$  ( $Z = 26$ ) attests to the enhanced stability of this isotope.

Figure 2 shows the abundance data for the elements up to Bi ( $Z = 83$ ), plotted vs. mass number to  $A = 210$ . Again, we see the regular decrease from  $A = 12$ , carbon, to the vicinity of  $A = 40$ , Ca and Sc, followed by the rise to mass 56, Fe. Then there is another smooth fall-off, out to the heaviest elements, with some localized peak structures apparent.

We can understand these trends, and even reproduce them in detail with our models of the reactions that occur in stars. These trends simply reflect the rules that science has uncovered about nuclei, about their reactions (nuclear dynamics) and their relative stabilities (nuclear structure).

The major characteristic of the data in Figure 2 is the general decrease in elemental abundance with increasing mass. This trend is a consequence of the nuclear reactions involved in elemental synthesis. Unusual stellar conditions, such as higher temperatures, are required to produce elements much heavier than helium, and different types of nuclear reactions come

into play as the energies go higher. In general, the reaction probabilities (called cross-sections) decrease, along with the numbers of available nuclei that can participate in a particular reaction, so it becomes increasingly difficult to produce ever heavier elements.

Properties that depend on the rules of nuclear stability are also evident in Figures 1 and 2. I will enumerate these rules below, but you can verify them yourself, simply by noting the known abundances of the stable isotopes of the elements, such as those listed in the chart of the nuclides, a small portion of which is shown in Figure 3.

Elements with even  $Z$  (atomic number) have many more stable isotopes than elements with odd  $Z$ . "Even-even" isotopes, which have even  $Z$  and even numbers of neutrons,  $N$ , and thus have even mass numbers,  $A$ , are more abundant than "even-odd" isotopes of the same  $Z$  (which have odd  $A$ ). Stable isotopes of odd- $Z$  elements usually have even  $N$  ("odd-even" nuclides); in fact, many odd- $Z$  elements have only one stable isotope. Stable "odd-odd" isotopes are rare. These characteristics account for the saw-tooth patterns, from even to odd  $Z$  and  $A$ , respectively shown in Figures 1 and 2.

In addition to these regularities, we know that certain regions of nuclei have enhanced stability; there are "closed shells" in nuclei just as there are in the electron orbitals in atoms. So-called "magic numbers"

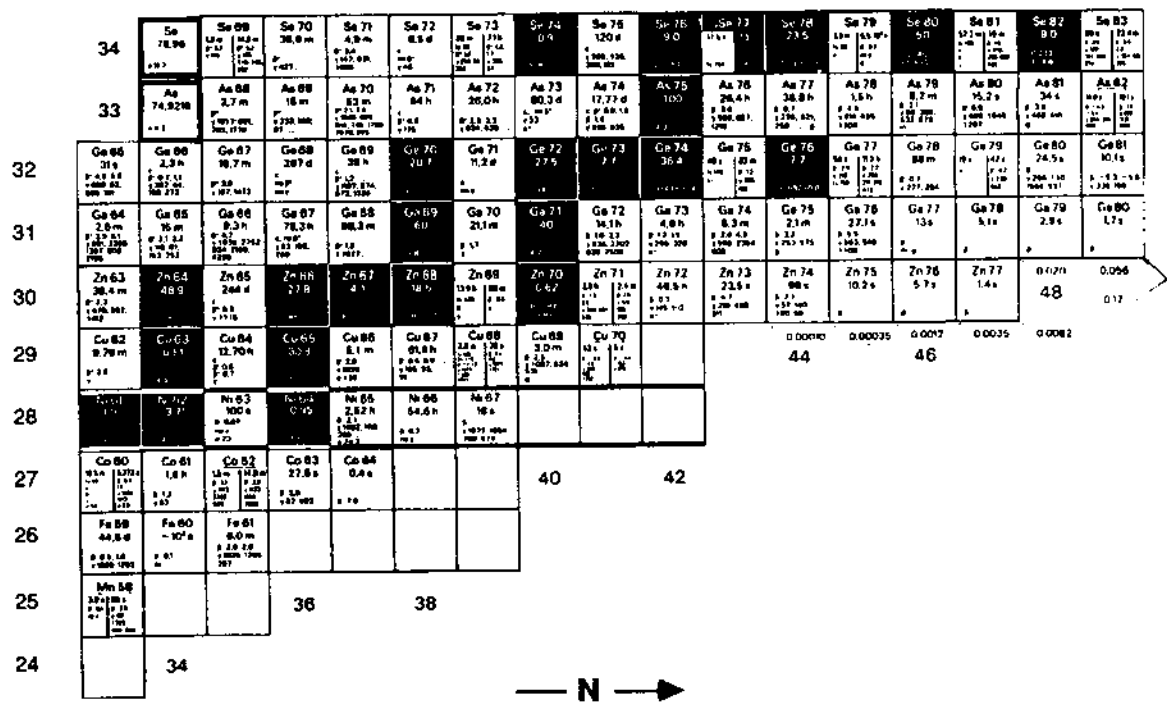
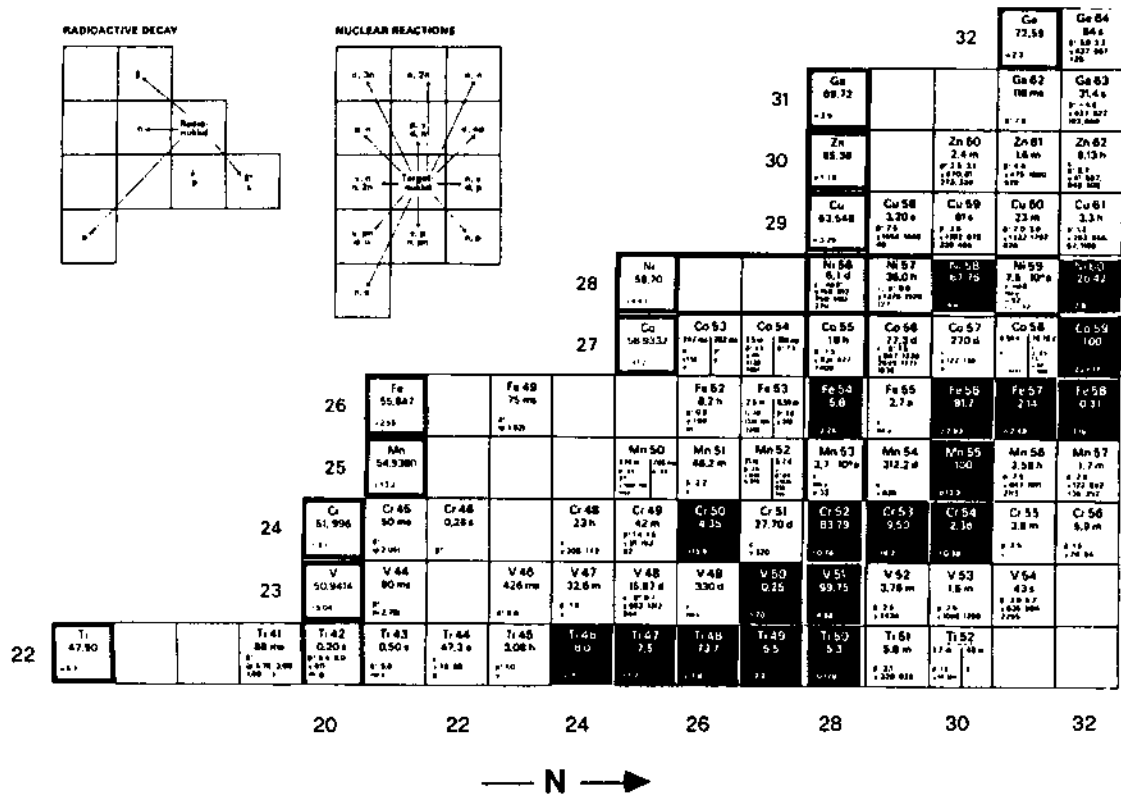


Figure 3. Portion of the Chart of the Nuclides (Knolls Atomic Power Laboratory, General Electric Co.).

identify these shell closures in nuclei, at  $Z$  and also  $N$  values of 2, 8, 20, 28, 50 and 82, and at  $N = 126$ . These stable regions are apparent in Figure 2, for example, in the peaks around  $A = 56, 130$  and  $208$ , which correspond, respectively, to isotopes around Fe and Ni (region of  $Z = 28$  and  $N = 28$ ); around Sn ( $Z = 50$  and  $A = 82$ ); and around Pb and Bi ( $Z = 82$  and  $N = 126$ ).

## V. NEUTRINOS AS PROBES OF THE STARS.

### Neutrino Production

As we have said, an important property of hydrogen burning, which accounts for energy production in our sun and other main sequence stars, is the emission of neutrinos. However, as we see in Table 1, where all of the pertinent reactions are listed, the pp reaction that we discussed does not account for all of the neutrinos produced in the sun. Other reactions are involved, which lead in part to  ${}^7\text{Be}$  and  ${}^8\text{B}$ . Figure 4 shows the calculated energy spectra of the different neutrino groups that are produced in the sun; intensities are given on the left-hand vertical scale. Note that more

than  $10^{11}$  pp neutrinos arrive at the Earth's surface per  $\text{cm}^2$  per second per MeV.

### Why Are We So Interested in Neutrinos?

Neutrinos are very unusual particles. They have essentially zero mass, and travel at the speed of light. They interact only via the weak nuclear force, so their probabilities of interacting with other forms of matter are extremely small. Whereas the cross-sections,  $S$ , for strong nuclear processes are about  $10^{-25} \text{ cm}^2$  (think of this number as an effective "size" or area of the nucleus), the cross-sections for neutrino interactions are much smaller,  $10^{-41} \text{ cm}^2$  or less. Because they are so extremely difficult to observe, neutrinos have been called "elusive," or "evanescent." Typical cross-section curves for neutrinos are also shown in Figure 4, for interactions with two different target atoms,  ${}^{71}\text{Ga}$  and  ${}^{37}\text{Cl}$  (right-hand vertical scale). More about these targets later.

These properties offer us a distinct advantage in studying the sun. Remember that hydrogen burning, the reactions listed in Table 1, takes place at the sun's

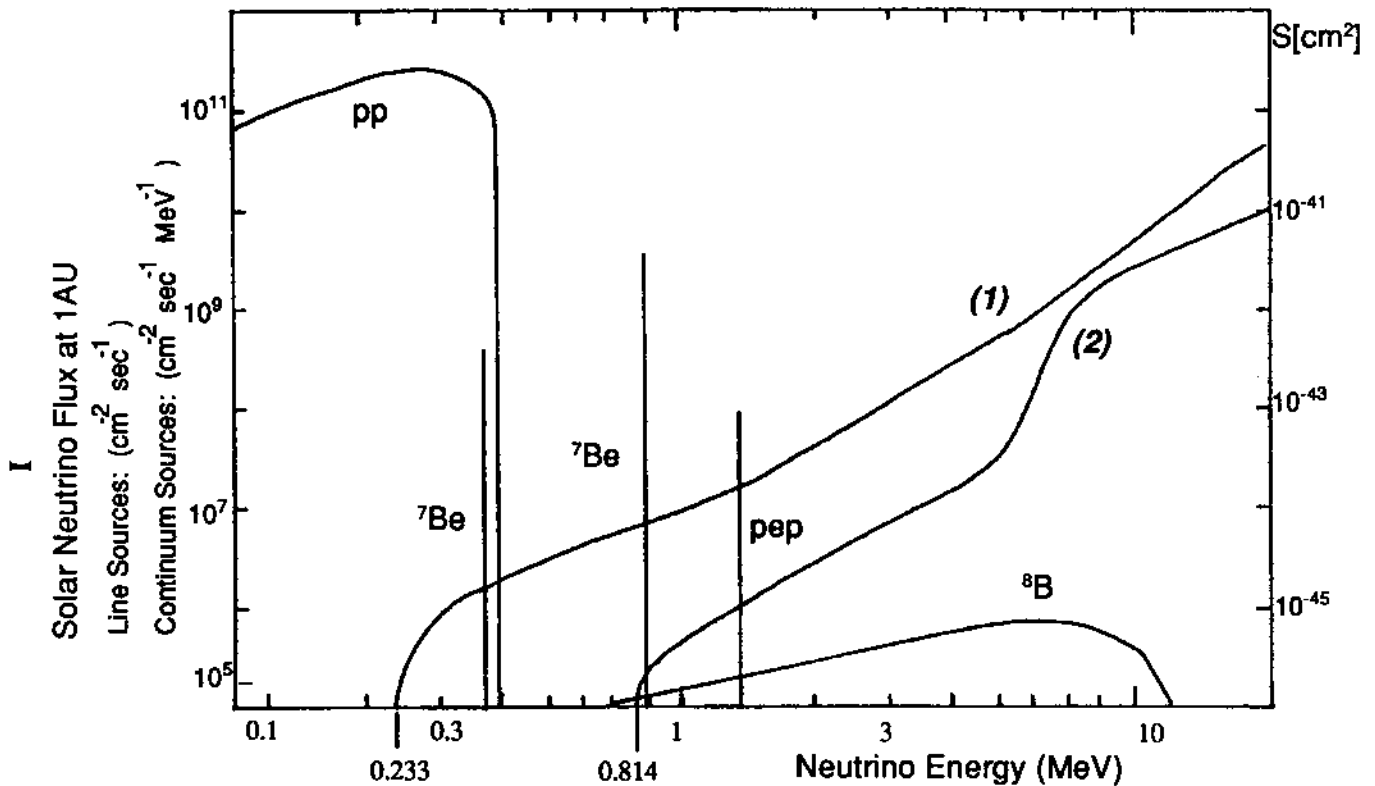


Figure 4. Neutrino energy spectra for solar hydrogen burning calculated with the standard solar model, with neutrino fluxes listed on the left-hand vertical axis. Cross-sections are also shown for neutrino capture in  ${}^{71}\text{Ga}$  (curve 1) and  ${}^{37}\text{Cl}$  (curve 2), with values listed on the right-hand vertical scale.

- (1)  ${}^{71}\text{Ga} (\nu_e, e) {}^{71}\text{Ge}$ ; 133 SNU  
55% p.p.  $\triangleright$   $R=I \cdot N \cdot S$   
theory
- (2)  ${}^{37}\text{Cl} (\nu_e, e) {}^{37}\text{Ar}$ ; 7.9 SNU  
76%  ${}^8\text{B}$
- 1 SNU  $\equiv 10^{-36}$   $\nu$  captures/sec-target atom

**Table 1.**  
**Neutrino Production in the Sun (Standard Solar Model)**  
**p-p Cycle.**

| Reaction.  | Branch.                                  | $\nu$ Type/Energy (MeV)            |
|--|--|------------------------------------|
| $p + p \rightarrow d + e^+ + \nu$  | 99.75% (p)                               | "pp": maximum = 0.42               |
| $p + e^- + p \rightarrow d + \nu$  | 0.25% (p)                                | "pep": 1.44                        |
| $d + p \rightarrow {}^3\text{He} + \text{gamma ray}$                     | 100% (d)                                 |                                    |
| ${}^3\text{He} + {}^3\text{He} \rightarrow {}^4\text{He} + p + p$        | 85% ( ${}^3\text{He}$ )                  |                                    |
| ${}^3\text{He} + {}^4\text{He} \rightarrow {}^7\text{Be} + \text{gamma}$ | 15% ( ${}^3\text{He}$ )                  |                                    |
| ${}^7\text{Be} + e^- \rightarrow {}^7\text{Li} + \nu$                    | 99.98% ( ${}^7\text{Be}$ )               | " ${}^7\text{Be}$ ": 0.38 and 0.86 |
| ${}^7\text{Li} + p \rightarrow {}^4\text{He} + {}^4\text{He}$            |  |                                    |
| ${}^7\text{Be} + p \rightarrow {}^8\text{B} + \text{gamma}$              | 0.02% ( ${}^7\text{Be}$ )                |                                    |
| ${}^8\text{B} \rightarrow {}^8\text{Be} + e^+ + \nu$                     |  | " ${}^8\text{B}$ ": maximum = 18.8 |
| ${}^8\text{Be} \rightarrow {}^4\text{He} + {}^4\text{He}$                |  |                                    |
| ${}^7\text{He} + p \rightarrow {}^4\text{He} + e^+ + \nu$                | $2 \times 10^{-5}\%$ ( ${}^7\text{He}$ ) | "hep": maximum = 18.8              |

core. Because of the high density of matter in the sun's interior, all of the products of these reactions, except the neutrinos, are effectively confined to that region because they undergo further interactions. The neutrinos can easily escape. Those that arrive at the Earth carry information about the nuclear processes that occur at the solar center. Thus, they can tell us how accurate are our theories of energy production in the sun and other stars.

#### How to Observe Neutrinos?

If neutrinos can escape the sun so easily, how can we expect to "see" them when they arrive in our laboratories? Well, their interaction cross-sections are very small but are not zero, so we can compensate for the small interaction probability by using a huge amount of detector material. To understand this statement, let's consider what happens when a neutrino interacts with a "detector." For the time being, we'll consider the interaction to be some vague form of "collision": the neutrino arrives, "hits" an atom in the detector, and causes some effect that we can observe. How can we describe this process in a general way?

We can think of a collision between a projectile and a target atom as depending on three key quantities: the number of projectiles that arrive per unit time ( $I$ ), the number of target atoms present in the detector ( $N$ ), and the cross-section for the collision to occur ( $S$ ). If we choose our units such that  $I =$  neutrinos per second per  $\text{cm}^2$ ,  $N =$  total number of target atoms, and  $S = \text{cm}^2$ , we find that the rate of collisions per second,  $R$ , is simply

$$R = I \times N \times S.$$

So what can we expect if we try to observe neutrinos? Let's take an example from Figure 4. At about 0.3 Mev,  $I = 10^{11}$   $\nu$  per  $\text{cm}^2$  per sec, and  $S = 10^{-45}$   $\text{cm}^2$  for  ${}^{71}\text{Ga}$ . If we want  $R$  to be about 1 interaction per second, then  $N$  must =  $10^{34}$  atoms of  ${}^{71}\text{Ga}$ . Since  $6 \times 10^{23}$  atoms of this isotope have a mass of 71 grams,  $N = 1.2 \times 10^{12}$  grams, or (since 1 ton = 1 thousand kilograms) 1.2 million metric tons! This quantity is enormous, especially since the world's annual production of gallium in the chemical industry is about 30 metric tons. What if we cut our expectations back to a more modest level, say  $R = 1$  per day? Then  $N$  becomes  $1.2 \times 10^6$  tons/86,400 seconds in a day, or 14 tons, still a huge amount, but manageable.

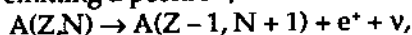
We have learned two important facts from these considerations: The amount of target required is huge, no matter what we do, and the expected event rate is very low, 1 per day.

#### Radiochemistry to the Rescue.

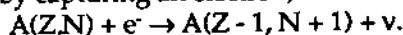
Now let's talk about how we can know that a neutrino has actually interacted with an atom in the detector. Table 1 shows us that each type of reaction in which a neutrino is produced in the sun is a form of beta decay: either a positron,  $e^+$ , is emitted, or an electron,  $e^-$ , is captured by the nucleus. Because of this relationship, the neutrinos emitted in these reactions are called "electron neutrinos" (there are two other types of neutrinos: one that is associated with the muon, and the second, with the tau particle). The electron neutrinos have an interesting property. Because they are emitted by nuclei in beta decay, they can also be captured by nuclei, initiating an "inverse beta decay."

For example, a radioactive nucleus with mass  $A$ ,

atomic number  $Z$  and neutron number  $N$ , can change one of its protons to a neutron in a decay process, by emitting a positron,



or by capturing an electron,



In some instances, the product nucleus,  $A(Z-1, N+1)$  is stable (not radioactive). In inverse beta decay, the neutrino is captured,  $A(Z-1, N+1) + \nu \rightarrow A(Z, N) + e^-$ . The target can be stable, and the product is radioactive.

So we have a handle on detecting the interaction of a neutrino. "All we have to do" is detect the presence of the radioactive product,  $A(Z, N)$ , in the stable target. No easy task when you remember that 1 radioactive atom per day is being produced in  $1.2 \times 10^{29}$  target atoms! The solution to this problem depends on chemistry. Because the product and the target have different atomic numbers, they have different chemical properties. As nuclear/radiochemists, we can devise specific, sensitive separations to isolate the small number of radioactive product atoms from the target, and can then detect them by observing their decay properties.

## VI. "THE SOLAR NEUTRINO PUZZLE."

### The $^{37}\text{Cl}$ Neutrino Detector.

Over twenty years ago, Dr. Ray Davis and colleagues at the Brookhaven National Laboratory built a radiochemical detector for solar neutrinos. They used  $^{37}\text{Cl}$  as the detector. When  $^{37}\text{Cl}$  captures a neutrino, it is transformed into radioactive  $^{37}\text{Ar}$ . The target used was 100,000 gallons (600 tons) of  $\text{C}_2\text{Cl}_4$ , perchloroethylene, the organic liquid that used in dry-cleaning establishments; the natural abundance of the isotope,  $^{37}\text{Cl}$ , in chlorine is 24%. To minimize backgrounds, the tank containing the Cl target was placed deep underground, in the Homestake gold mine in Lead, South Dakota. The 5,000 feet of rock above the detector served as a shield, to block out cosmic rays, reduce the levels of naturally occurring radioactivity such as U and Th, etc., but did not cut out any of the neutrinos.

Because Ar is a noble gas, the radioactive product can be easily removed from the liquid target in a stream of gas, and the decay of the  $^{37}\text{Ar}$  can then be detected in a low-level, gas-filled proportional counter. Davis' experiment has been running for more than two decades. His experimental result, averaged over that time period, for the neutrino-induced production of  $^{37}\text{Ar}$  is 2.1 SNU, where one Solar Neutrino Unit equals  $10^{-36}$  neutrino captures per target atom per second. The "problem" with this result is that the standard solar model predicts a value of 7.9 SNU. The discrepancy between the experiment and the theory is a factor of 4, well beyond the statistical errors in either number. These missing neutrinos have come to be known as "The Solar Neutrino Puzzle."

### How to Explain the Discrepancy?

The result from the  $^{37}\text{Cl}$  is a measurement of the quantity  $R = I \times N \times S$ . What might be the cause of the discrepancy?

First of all, is the measured value,  $R$ , correct? The consensus has developed in the scientific community that Davis' experimental result is trustworthy. No serious experimental errors exist that would explain away the discrepancy.

$N$  is known accurately since it depends only on the weight and composition of the target material. What about  $S$ ? It is thought that the cross-sections for neutrino capture in  $^{37}\text{Cl}$ , which come from experimental data on the radioactive decay of  $^{37}\text{Ar}$  and from nuclear reaction studies, are known accurately enough.

The only questionable quantity remaining is  $I$ , the flux of neutrinos arriving from the sun. A few intriguing possibilities remain.

One is that perhaps we don't know the sun and its nuclear processes as well as we think we do. For example, if you look at the cross-section vs. energy curve for  $^{37}\text{Cl}$  in figure 4, you'll see that it goes to zero at 814 keV; this is the "threshold" of the  $^{37}\text{Cl}$  detector for neutrinos. So,  $^{37}\text{Cl}$  is primarily sensitive (76%) to the " $^{\text{B}}$ " neutrinos from the sun. Notice that the neutrinos from this branch have the lowest intensities and very high energies. As a result, their production rate in the sun is very sensitive to the value of the sun's central temperature, as well as to other properties of the solar core. However, theoretical models that have taken variations of these parameters into account have as yet been unable to explain the discrepancy.

Another very exciting possibility is that the sun is behaving "properly," but that the neutrino is "doing wild and wonderful things." Perhaps the "correct" number of neutrinos is being manufactured in the sun, but some of them are being "waylaid" on their trip to the Earth. I'll have more to say about this possibility at the end.

### $^{71}\text{Ga}$ , A New Kind of Neutrino Detector.

If you wanted to verify Davis' result, you might think of repeating his experiment, and possibly increasing the sensitivity by greatly increasing the amount of Cl used. However, a more attractive approach is to create a new kind of neutrino detector, with properties that are different from the Cl detector, so you can approach the neutrino problem "from a new angle."

Several schemes for new radiochemical neutrino detectors have been devised, but most have never left the drawing board. One that has is the  $^{71}\text{Ga}$  detector. Figure 4 shows the advantage of this detector: note that the cross-section curve for  $^{71}\text{Ga}$  goes all the way down to a threshold of 233 keV, much below that of  $^{37}\text{Cl}$ . So  $^{71}\text{Ga}$  is sensitive to all of the different solar neutrino branches, including a large contribution (55%) from the primary "pp" branch. Thus, the result from a  $^{71}\text{Ga}$



detector should be fairly insensitive to conditions at the solar core, a great advantage in the experiment.

Because of its low threshold, the predicted capture rate in  $^{71}\text{Ga}$  is 133 SNU, much larger than for  $^{37}\text{Cl}$ . So an event rate of 1 per day can be obtained with only 30 tons of gallium (a good thing since 30 tons of Ga cost 15 million dollars!).

When  $^{71}\text{Ga}$  captures a neutrino, it forms radioactive  $^{71}\text{Ge}$ . It so happens that Ge forms a volatile chloride,  $\text{GeCl}_4$ , so that the  $^{71}\text{Ge}$  can be swept out of the Ga target in a stream of gas, analogously to what is done in removing  $^{37}\text{Ar}$  from Cl. Ge also forms a gaseous hydride,  $\text{GeH}_4$ , that is quite suitable for gas-proportional counting. In many ways, the Ga experiment is similar to the Cl one. The important difference is that a very different range of neutrino energies can be probed with Ga, all the way down to the all-important pp branch, which starts the hydrogen burning chain in the sun.

At present, not one but two  $^{71}\text{Ga}$  detectors are beginning operation. The two experiments are similar in many ways, but they differ in some significant aspects of their radiochemistry. One detector, which employs an aqueous solution of  $\text{GaCl}_3 + \text{HCl}$ , is nearing completion in a tunnel under the Appennine Mountains in Italy, at the Gran Sasso National Laboratory. My group at Brookhaven National Lab is a participant in this experiment; it is called GALLEX, a consortium of 11 laboratories, mainly in Europe. The other experiment, which uses Ga metal, a liquid at  $30^\circ\text{C}$ , has started to operate in a tunnel under the Caucasus Mountains in the Soviet Union. The experiment is called SAGE, a collaboration between laboratories in the U.S.S.R. and the U.S.A. As you might expect, the chemistry of removing volatile  $\text{GeCl}_4$  from molten Ga metal is more involved than removing it from aqueous Ga solution.

What can we expect from these Ga detectors? First of all, the two Ga experiments had better get the same answer! More importantly, they can tell us if the flux of low-energy neutrinos arriving here from the sun, especially from the pp branch, is considerably less than that predicted by the standard solar model, in accord with the measured low result from the  $^{37}\text{Cl}$  experiment for the  $^8\text{B}$  branch.

## VII. NEW NEUTRINO PROPERTIES??

### Does the Neutrino Have Mass?

An idea has been developed in elementary particle physics in recent years that is relevant to the solar neutrino puzzle. Simply stated, the idea is that the three different types or "flavors" of neutrinos (those associated with the three leptons, the electron, the muon, and the tauon) are all members of one "family." As such, they can "change places" in the family; e.g., an electron neutrino could fairly easily be transformed into a muon neutrino. The importance of this transformation for our discussion is that only electron neutrinos can initiate the inverse beta-decay process. If an electron

neutrino produced in the sun is changed into a muon neutrino (or even an electron antineutrino) during its journey to the Earth, the neutrino will be "sterile" when it arrives at the radiochemical neutrino detector; it will be incapable of converting  $^{37}\text{Cl}$  to  $^{37}\text{Ar}$ , or  $^{71}\text{Ga}$  to  $^{71}\text{Ge}$ .

The significance of this notion lies in the fact that the only way that the three neutrinos can form such a family is if they have different masses. These masses would be small, of the order of  $mc^2 =$  a few eV or even less, whereas the mass of the electron is 511 eV. Nevertheless, a neutrino that has mass would represent a major change in our thinking about this particle.

If this eventuality turns out to be the answer to the solar neutrino puzzle, we will be able to say that our study of the largest body in our solar system, the sun, told us something profound about the smallest body in the system of elementary particles, the neutrino.

## Suggested Reading

### General.

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